INFORMATION TO USERS

This manuscript has been reproduced from the microfilm master. UMI films the text directly from the original or copy submitted. Thus, some thesis and dissertation copies are in typewriter face, while others may be from any type of computer printer.

The quality of this reproduction is dependent upon the quality of the copy submitted. Broken or indistinct print, colored or poor quality illustrations and photographs, print bleedthrough, substandard margins, and improper alignment can adversely affect reproduction.

In the unlikely event that the author did not send UMI a complete manuscript and there are missing pages, these will be noted. Also, if unauthorized copyright material had to be removed, a note will indicate the deletion.

Oversize materials (e.g., maps, drawings, charts) are reproduced by sectioning the original, beginning at the upper left-hand corner and continuing from left to right in equal sections with small overlaps. Each original is also photographed in one exposure and is included in reduced form at the back of the book.

Photographs included in the original manuscript have been reproduced xerographically in this copy. Higher quality 6" x 9" black and white photographic prints are available for any photographs or illustrations appearing in this copy for an additional charge. Contact UMI directly to order.

University Microfilms International
A Bell & Howell Information Company
300 North Zeeb Road, Ann Arbor, MI 48106-1346 USA
313/761-4700 800/521-0600

Order Number 9416060

Stellar oxygen abundances

King, Jeremy R., Ph.D.

University of Hawaii, 1993

U·M·I

300 N. Zeeb Rd.
Ann Arbor, MI 48106

STELLAR OXYGEN ABUNDANCES

A DISSERTATION SUBMITTED TO THE GRADUATE DIVISION OF THE
UNIVERSITY OF HAWAII IN PARTIAL FULFILLMENT OF THE
REQUIREMENTS FOR THE DEGREE OF

DOCTOR OF PHILOSOPHY
IN
ASTRONOMY

DECEMBER 1993

By

Jeremy R. King

Dissertation Committee:

Ann M. Boesgaard, Chairperson

James N. Heasley

Barry J. Labonte

Alan N. Stockton

Lawrence J. Wallen

© Copyright 1993

by

Jeremy R. King

iii

Acknowledgements

An undertaking such as this is not possible without the help and support of many people. First and foremost of these was my advisor Dr. Ann Boesgaard. Much of what I have learned about stellar spectroscopy over the past five years was from her. The responsibilities she assigned me and the trust she placed in me during this time were valued and her patient and friendly manner were much appreciated. When I originally approached Ann about doing an observational thesis on oxygen, she responded enthusiastically and included expanded support for this work in her proposal to NSF; I am grateful to her for giving me such freedom in the choice of a thesis topic. Additionally, Ann gratefully provided me with brief respites from my research by allowing me to look after her house, which has such a lovely and peaceful ocean view. I also much enjoyed tending to Daiquiri, her “astronomical cat”, whose persistent and best efforts could not permanently keep me on an early rising schedule. Many thanks and much aloha, Ann, for your encouragement, genuine interest, and the opportunity to work with you.

Dr. James Heasley is the person most responsible for me coming (and, ironically, almost not coming) to Hawai‘i. He has given me much guidance while Ann Boesgaard was a visiting scientist at CFHT and Eberly chair at Penn State. Although it is not completely clear to me, I have probably benefited from the many times he has grumbled at me, my thesis, and my TeX skills. The idea to observe oxygen in open clusters was one he pushed strongly for and originally suggested.

Many thanks to Dr. Barry Labonte for his humor, his valued input, and his broad knowledge of stellar astronomy. It was Barry’s idea to use the 400 mm camera on the FOS to perform most of the stellar oxygen observations when I was seemingly out of options. Barry also accompanied me to Mees Solar Observatory and carried out the lion’s share of the solar observations there. Dr. Alan Stockton made the stellar observations a reality by designing and directing the construction of the camera mount and the appropriate slit. He also contributed financially to that effort and I am very grateful to him. Dr. Larry Wallen graciously accepted the task of being the outside member on my committee.

For the past two years, I have also been fortunate enough to work with Dr. Constantine Deliyannis. I appreciate his enthusiastic interest in my thesis work and his constant encouragement. He has given me useful feedback and detailed comments on practically all the material in this dissertation. I have certainly benefited from discussions with him concerning stellar models, primordial Li, and cosmology. He also kindly provided one of the isochrones shown in Chapter 3. I also thank him for his undying patience when my attention could not be focussed on our collaborative projects unrelated to my thesis. Additionally, I am grateful for his time and effort in writing job recommendations for

me. Most of all, I have valued his friendship in offering me counsel, advice, and opinion. His and his wife Deborah's hospitality on several occasions have provided needed and enjoyable diversions from ordinary graduate student life. Also I should like to thank him for an enjoyable observing run in Chile and for preserving my life via his nimble (harrowing?) driving through the deathtraps that are the streets of Santiago.

I have been financially supported the past three years by NSF grant AST-9016778 to Dr. Ann Boesgaard. CCD astronomy at the UH 2.2-m is supported by NSF grant AST-9020680. Mees Solar Observatory is supported by NSF grant AST-9115038 and NASA grant NAGW 1542. Additionally, it is a genuine pleasure to acknowledge the gracious financial support of the Honolulu chapter of the ARCS Foundation, Inc. during my first four years at UH. Mss. Diane Tokumura and Georgette Albert are to be thanked for insulating me, to the extent possible, from the usual realms of senseless, onerous bureaucracy and getting me to where I had to go. The UH 2.2-m telescope operators, Mssrs. Dave Woodworth and Frank Cheigh, skillfully and repeatedly placed my objects on a 0.7" slit and offered much additional assistance. Dr. Andrew Pickles, the UH 2.2-m day crew, and the IfA machine shop also played important roles in making the UH 2.2-m observations discussed here possible. Dr. Richard Canfield allowed use of Mees Solar Observatory to obtain the solar observations discussed in Chapter 5; additionally, I am grateful to him for providing travel support to obtain the observations. Ms. Kristin Blais helped me utilize the IDL software package in the reduction of the solar data and wrote several useful codes for my benefit. She also bore the frustrating burden of trying to explain spatially scanning spectroscopy to a point-source spectroscopist. During my dissertation research, I have made extensive use of the SIMBAD database, operated at CDS, Strasbourg, France.

I am grateful to the Kitt Peak National Observatory TAC for granting me time to perform the Hydra multi-object spectrograph observations. Full support of travel and observing expenses by KPNO is gratefully acknowledged. Dr. Sam Barden answered many questions about Hydra prior to the run and his guidance during the early observations is appreciated. The KPNO observations would not have gone as smoothly as they did without the competent assistance of the LTO, Ms. Jennifer Hedden. Dr. Jeanette Barnes spent the good part of a day in Tucson leading me through the Hydra reductions. She and Dr. Caty Pilachowski also have answered numerous questions over the past few years concerning the reduction of spectra using IRAF. Dr. Kenneth Janes, who initially kindled my interest in stellar astronomy, kindly provided a CCD image of NGC 188 which allowed me to obtain the astrometry for the Hydra observations. He also provided me with his unpublished photometry of NGC 188.

I am extremely grateful to Dr. Robert Kurucz, who unselfishly provided me with his unpublished model atmospheres, colors, and Fe log gf values, with which much of the analysis presented here was conducted. Thanks are also due Dr. Monique Spite and Dr. Lewis Hobbs for providing me with several of their halo star spectra. Dr. Spite also allowed the use of and provided guidance with her LTE abundance software package, RAI10. Dr. Michael Bessell kindly provided me with several pages of his personal notes concerning the reanalysis of O abundances in giants. Dr. Eric Feigelson made his SLOPES software package available to me to perform the regressions discussed in Chapter 4. Dr. Pierre Demarque made his isochrones for NGC 188 available for my use in Chapter 6. Dr. Jocelyn Tomkin provided me with his LTE and NLTE solar calculations for the O I triplet.

Finally, I must express my appreciation to my family and friends. I have managed to maintain my (in)sanity over the past five years with the support, concern, and valued friendship of Kristin Blais, Dave Kane, Niki Kapsambelis, Nancy Kohrs, Emma Murray, Mark Nadeau, and Ron and Diane Nadeau. I wish to thank my extended family for their interest and support, which were valued more than they probably knew. Special thanks go to my sister Kathy. Additionally, my delightful nieces Janna and Jaime and my nephew Nathan have been a source of enduring inspiration as well as pride to me. Lastly, I should like to dedicate this dissertation to my parents, Philip and Millideen, who were (and are) the most influential teachers in my life. Their wisdom, encouragement, interest, and love have been immeasurably important to me. Many thanks to one and all.

Abstract

We address a variety of issues concerning stellar oxygen abundances. We first investigate the discrepancy in O abundances of halo stars as derived from the 7774 Å O I triplet and other lines and propose a possible resolution to this discrepancy. The 7774 Å O I triplet equivalent widths of Abia & Rebolo [1989, AJ, 347, 186] are found to be systematically too high. It is also argued (from a number of lines of evidence) that present effective temperature values used in halo star abundance studies may be ~ 150 K too low. New color- T_{eff} relations are derived for metal-poor stars. Using the T_{eff} values from these relations, improved equivalent widths for the 7774 Å O I triplet, and revised $[\text{Fe}/\text{H}]$ values (consistent with our revised T_{eff} values), we redetermine $[\text{O}/\text{Fe}]$ ratios for a handful of metal-poor stars. The mean ratio is $[\text{O}/\text{Fe}] = +0.52$ and no dependence on T_{eff} or $[\text{Fe}/\text{H}]$ is found for these stars. Next, possible cosmological implications of the hotter T_{eff} scale are discussed along with additional evidence supporting the need for a higher temperature scale for metal-poor stars. Based on Li abundances in the literature, our T_{eff} scale leads to a Spite plateau value of $N(\text{Li}) = 2.28 \pm 0.09$ (one standard deviation). A conservative minimal primordial value of $N(\text{Li}) = 2.35$ is inferred. While agreement with standard big bang nucleosynthesis models is not particularly good, if errors in the observations and models are considered, consistency with the model is still achieved in the range $0.012 \leq \Omega_b \leq 0.017$ (for $H_0 = 100$ km/s/Mpc). The revised T_{eff} scale raises the observed B/Be ratio of HD 140283 from 10 to 12, making its value more comfortably consistent with the production of the observed B and Be by ordinary spallation. Our T_{eff} values are found to be in good agreement with values predicted from both the Victoria and Yale isochrone color- T_{eff} relations. Thus, it appears likely that no changes in globular cluster ages would result (at least for those clusters having $[\text{Fe}/\text{H}] \sim -1.3$). Next, we examine the location of the break in the $[\text{O}/\text{Fe}]$ *versus* $[\text{Fe}/\text{H}]$ plane in a quantitative fashion for the first time. Analysis of a relatively homogeneous data set does not favor any unique break point in the range $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$. Further analysis suggests, however, that the break is more likely to occur near $[\text{Fe}/\text{H}] \sim -1.0$ rather than recently claimed values of $[\text{Fe}/\text{H}] \sim -1.7$. Additionally, the red giant O dataset used in this analysis also indicates that $[\text{O}/\text{Fe}]$ is constant in the halo (at least for $[\text{Fe}/\text{H}] \geq -3$), in agreement with the results for halo dwarfs. We qualitatively discuss the implications of our results with particular attention to the possibility of a delay between the end of halo formation and the beginning of star formation in the disk. We find that the gap in the observed $[\text{O}/\text{H}]$ distribution, noted by Wheeler *et al.* [1989, ARAA, 27, 279], persists despite the addition of more O data and *may* betray the occurrence of such a hiatus in star formation. It is noted that the slope of the $[\text{O}/\text{Fe}]$ *versus* $[\text{Fe}/\text{H}]$ relation for $[\text{Fe}/\text{H}] \geq -1$ depends on the statistical regression utilized. Given this sensitivity, alleged “observed” $[\text{O}/\text{H}]$ –age

relations, which do not use truly observed O abundances (but, rather, adopt O abundances based on Fe abundances), should be regarded with caution. Additionally, the implications of these “observed” relations on issues regarding Galactic chemical evolution should also be regarded with caution. Next, systematic effects on O abundances derived from the 6300 Å [O I] and 7774 Å O I lines are considered. New Solar observations presented here confirm the disagreement between the observed 7774 Å O I equivalent widths and LTE model calculations at low μ . However, we stress that Solar O abundance determinations made from *flux* spectra are in very good agreement with the meteoritic value. It is noted that the 6300 Å [O I] equivalent width value appears to be uncertain for the Sun. Given this uncertainty, the inability of authors to reproduce each others’ 6300 Å O abundances, and the results of recent quasi-two-stream calculations, we do *not* believe it can be claimed *a priori*, as is usually done, that these abundances are more reliable than those derived from the 7774 Å O I triplet. We empirically assess the situation by deriving abundances for a selection of relatively metal-rich F and G dwarfs. For $T_{\text{eff}} \geq 6200 - 6300$ K, the 7774 Å O I abundances are substantially larger than the 6300 Å [O I] abundances; this is interpreted as the onset of observable NLTE effects seen in early-type stars. For lower temperatures, however, we find no systematic difference between the 6300 and 7774 Å abundances, a result in conflict with that of Nissen & Edvardsson [1992, A&A, 261, 255]. It seems likely that the discrepancy resides in the model atmospheres used in the analysis. In any case, if our hotter T_{eff} scale for metal-poor stars is correct, then it is noted again that there appears to be no discrepancy between the 6300 Å and 7774 Å O abundances of halo stars. Thus, very large effects of NLTE conditions or atmospheric thermal inhomogeneities are ruled out; however, small effects (*i.e.*, ~ 0.05 dex) due to NLTE departures or thermal inhomogeneities cannot be excluded. Finally, O abundances are determined in 8 open clusters or moving groups. Poor agreement is found with the recent determinations Garcia Lopez *et al.* [1993, ApJ, 412, 173] for the Hyades and UMa group. We discuss why we believe our abundances to be more accurate. A very clear relation between cluster age and O abundance is seen; this is in stark contrast to the lack of any relation between age and Fe abundance in the same clusters. It is argued that, despite possible *a priori* objections, O abundances may prove to be a superior chronometer (as others have suggested) in the study of Galactic chemical evolution. The behavior of Mg being overabundant with respect to iron for the older clusters as noted by Friel & Janes [1993, A&A, 267, 75] is not seen for O. Rather, the opposite effect is seen—our [O/Fe] ratios appear to be larger for the younger clusters. The O abundances in the younger clusters are super-solar and are significantly larger than those seen in H II regions, planetary nebulae, and supergiants. It is suggested, and supported with observational evidence, that this may be due to incomplete stellar models, which do not consider phenomena such as rotational mixing and the fact

that a significant fraction of O in gaseous nebulae is locked up in dust grains. With this in mind, we reinterpret existing planetary nebulae and supergiant abundances as being in good agreement with our cluster results. If this is correct, then existing inferences made about extragalactic O abundances or O abundance gradients in the Galaxy based on analysis of planetary nebulae or nebulae may be incorrect. Incomplete stellar models may also imply that O abundances derived in some halo giants may not reflect initial O abundances that these stars formed with. Again, we reinterpret existing red giant O data as indicating that $[O/Fe] \sim +0.5$ in the halo. Investigation of rotational models having a range of metallicities and masses is urgently needed to shed further light on these important topics. Examining our results in a broader sense, we suggest that: a) the formation of the Galactic halo was a slow process but b) did not involve the merger of independent “fragments” c) Type Ia supernovae are dominated by CO-He white dwarf systems having Fe production timescales of a few Gyr d) The hiatus between the end of halo formation and the beginning of star formation in the disk, possibly required if CO-He white dwarfs are the dominant source of Type Ia supernovae, may be confirmed as the observed gap in the $[O/H]$ distribution e) The long timescale ($10^9 - 10^{10}$ yr) of Fe production by CO-He white dwarf systems is also seen to be consistent with the lack of any correlation between age and $[Fe/H]$ in our open cluster sample and the large scatter present in the $[Fe/H]$ *vs.* age relations for field stars. These latter conclusions are not firm, however, since there are other suggestions from the data which seem to be consistent with fragment merger models. Observations to determine $[O/Fe]$ ratios in bulge stars might shed some additional light on the subject.

Table of Contents

Acknowledgements	iv
Abstract	vii
List of Tables	xi
List of Figures	xii
Preface	xiv
Chapter 1. Introduction	1
Chapter 2. A Resolution to the 7774 Å O I Abundance Discrepancy	3
Chapter 3. Possible Cosmological Implications of a Hotter Temperature Scale for Halo Stars	37
Chapter 4. The Location of the Break in the [O/Fe] <i>vs.</i> [Fe/H] Plane	59
Chapter 5. Systematic Effects on O Abundances from the 6300 Å [O I] and 7774 Å O I Lines	81
Chapter 6. Oxygen Abundances in Open Clusters	117
Chapter 7. Conclusions, Summary Comments, and Future Work	158
Literature Cited	165

List of Tables

1. Comparison of 7774 Å O I Triplet Equivalent Widths	21
2. Photometric Data for Metal-Poor Stars	22
3. Spectroscopic T_{eff} Values from Line Ratios	23
4. Comparison of Spectroscopic and Photometric Temperatures I.	24
5. Comparison of Spectroscopic and Photometric Temperatures II.	25
6. Reanalysis of [O/Fe] Ratios in Metal-Poor Stars	26
7. Rederived Halo Star ^7Li Abundances	53
8. Halo Star T_{eff} Comparison	54
9. Solar O I Triplet Equivalent Widths	102
10. Comparison of Solar 7774 Å O I Equivalent Widths	103
11. Stellar Parameters	104
12. Equivalent Widths and Abundances	106
13. Pleiades Parameters and Abundances	142
14. Coma Parameters and Abundances	143
15. UMa Moving Group Parameters and Abundances	144
16. Praesepe Parameters and Abundances	145
17. Hyades Parameters and Abundances	146
18. NGC 752 Parameters and Abundances	147
19. M 67 Parameters and Abundances	148
20. NGC 188 Parameters and Abundances	149
21. Cluster Abundances and Parameters	150

List of Figures

Figure 1. Kurucz (1991) Model $(b - y) - T_{\text{eff}}$ Relations	27
Figure 2. Comparison of $(b - y) - T_{\text{eff}}$ Relations	28
Figure 3. Kurucz (1991) Model $(B - V) - T_{\text{eff}}$ Relations	29
Figure 4. Comparison of $(B - V) - T_{\text{eff}}$ Relations	30
Figure 5. Kurucz (1991) Model $(V - K) - T_{\text{eff}}$ Relations	31
Figure 6. $T(b - y) - T(B - V)$ vs. $[\text{Fe}/\text{H}]$ for Magain (1987) Relations	32
Figure 7. $T(b - y) - T(V - K)$ vs. $[\text{Fe}/\text{H}]$ for Magain (1987) Relations	33
Figure 8. $(V - K) - T_{\text{eff}}$ Relation for Pleiads	34
Figure 9. $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ for Tomkin <i>et al.</i> (1989) Subsample	35
Figure 10. $[\text{O}/\text{Fe}]$ vs. T_{eff} for Tomkin <i>et al.</i> (1989) Subsample	36
Figure 11. $[\text{Fe}/\text{H}]$ vs. Line Excitation Potential for HD 89499	55
Figure 12. Primordial Nucleosynthesis Model Predictions	56
Figure 13. Comparison of Bergbusch & Vandenberg (1992) Isochrone with M92 Fiducial	57
Figure 14. Comparison of Chaboyer <i>et al.</i> (1992) Isochrone with M92 Fiducial	58
Figure 15. Combined Red Giant and Dwarf Sample in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ Plane	74
Figure 16. A Sample Regression in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ Plane	75
Figure 17. Combined ^9Be Sample in the $\log n(\text{Be})/n(\text{H})$ vs. $[\text{Fe}/\text{H}]$ Plane	76
Figure 18. Polynomial and Spline Regressions in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ Plane	77
Figure 19. Combined Red Giant and Dwarf Sample in the $[\text{Fe}/\text{O}]$ vs. $[\text{O}/\text{H}]$ Plane	78
Figure 20. Combined Red Giant and Dwarf Sample in the $[\text{Fe}/\text{H}]$ vs. $[\text{O}/\text{H}]$ Plane	79
Figure 21. $[\text{Fe}/\text{H}]$ Distribution of Combined Red Giant and Dwarf Sample	80
Figure 22. Sample Spectra of the Solar 7774 Å O I Triplet	108
Figure 23. Predicted and Observed Solar 7774 Å Triplet Equivalent Widths vs. μ	109
Figure 24. Observed 7774 Å Equivalent Widths vs. Predicted NLTE Values	110
Figure 25. Spectrum of the 7774 Å Region for HR 8969	111
Figure 26. $\Delta \log g$ vs. Kurucz (1992) $\log g$ Value	112
Figure 27. 6300 Å [O I] Abundances vs. 7774 Å O I Abundances	113
Figure 28. Difference of 7774 Å O I and 6300 Å [O I] Abundances vs. T_{eff}	114
Figure 29. $\Delta[\text{O}/\text{H}]$ from Nissen & Edvardsson (1992) vs. Their T_{eff} Values	115
Figure 30. Difference of 7774 Å O I and 6300 Å [O I] Abundances vs. $[\text{Fe}/\text{H}]$	116

Figure 31. Sample 7774 Å Region Spectra of Open Cluster Stars	151
Figure 32. [Fe/H] <i>vs.</i> Age for the Open Cluster Sample	152
Figure 33. [O/H] <i>vs.</i> Age for the Open Cluster Sample	153
Figure 34. [O/H] <i>vs.</i> [Fe/H] for the Open Cluster Sample	154
Figure 35. [O/Fe] <i>vs.</i> Age for the Open Cluster Sample	155
Figure 36. [O/Fe] <i>vs.</i> [Fe/H] for the Open Cluster Sample	156
Figure 37. [Fe/O] <i>vs.</i> [O/H] for the Open Cluster Sample	157

Preface

Here, I alert the reader to the publication status and plans for this dissertation as of this date. Chapters 1 and 2 have been published, practically verbatim, in the September 1993 volume of *The Astronomical Journal* (AJ, 106, 1206). Chapter 4 is in press and is scheduled for publication in the January 1994 volume of *The Astronomical Journal*. Chapter 3 is also in press and is scheduled for publication in the March 1994 volume of *The Astronomical Journal*. Presently, it is planned that Chapters 5 and 6, the former a joint publication by Dr. Ann Boesgaard and myself, will be submitted for publication in *The Astronomical Journal* in early 1994 so that publication might occur in mid-to-late 1994. Just as for Chapters 1 and 2, we note that the material unpublished as of this date is likely to appear essentially verbatim, though at least slight cosmetic changes and addition of supplementary material will certainly occur, in the journal articles.

Finally, it should be mentioned that I have obtained much data as part of my dissertation observations that are not presented here. This is due to sensible length constraints (*i.e.*, brevity), time constraints (due to uncooperative weather, most of the data obtained was during the last year, resulting in a “glut” of sorts), and other practical constraints (*e.g.*, trying to narrow the focus of the dissertation and being able to graduate so I can achieve gainful employment, *etc.*). Current plans call for at least three more papers in the “Stellar Oxygen Abundance” series in *The Astronomical Journal* over and above those mentioned in the first paragraph. These papers will present observations and discuss the: 1) O abundances and [O/Fe] ratio in halo stars 2) O abundances in a potential sample of (so-called) “thick disk” stars and 3) the age-[O/H] relation among solar-neighborhood field stars. The first paper is likely to be another collaborative effort with Dr. Ann Boesgaard. We hope to be able to present our results on these three topics in *The Astronomical Journal* within the next three years or so.

Chapter 1. Introduction

1. Overview

Oxygen is the most abundant metal in the Galaxy. As such, O abundances provide important constraints on our understanding of Galactic chemical evolution (*e.g.*, Wheeler, Sneden & Truran 1989) and Galaxy formation (*e.g.*, Gilmore, Wyse & Kuijken 1989; Matteucci & Francois 1992). The run of $[\text{Fe}/\text{O}]$ *vs.* $[\text{O}/\text{H}]$ bears on such issues as the nature and temporal variability of the star formation rate and initial mass function in the Galaxy (Wheeler *et al.* 1989). The value of the $[\text{O}/\text{Fe}]$ ratio and its behavior with respect to $[\text{Fe}/\text{H}]$ in low metallicity stars has important implications for the possible “G dwarf-problem” in the Galactic halo (Ryan & Norris 1991a). The $[\text{O}/\text{Fe}]$ ratio is also an important parameter in models of Galactic chemical evolution constructed for comparison with observed $[\text{Be}/\text{H}]$ ratios in attempts to constrain cosmological models of primordial nucleosynthesis (*e.g.*, Ryan *et al.* 1992). Gilmore *et al.* (1989) assert that the $[\text{O}/\text{Fe}]$ ratio is the most important observational constraint on the timescale of Galactic formation. Indeed, the observed ratio of $[\text{O}/\text{Fe}]$ in the halo has been used to constrain the duration of halo collapse via comparison with models (Matteucci & Francois 1992, Smecker-Hane & Wyse 1992). Precise relative values of the $[\text{O}/\text{Fe}]$ ratio (and the observational uncertainties) are important for the evaluation of models of Galaxy formation (*e.g.*, Searle & Zinn 1978) due to mergers of independent fragments (Gilmore *et al.* 1989). Knowledge of the $[\text{O}/\text{Fe}]$ ratio in the halo is also crucial for making more accurate absolute and relative age determinations for globular clusters via comparison with theoretical isochrones (VandenBerg 1988, VandenBerg & Stetson 1991).

2. Observing Oxygen in Stellar Spectra

The element traditionally used in the study of issues related to those above is iron. Lines of Fe in stellar spectra are numerous and easily observed. However, this is not the case for O. Useful permitted lines of neutral O (the dominant ionization species in solar temperature stars) exist in the near-IR region (*e.g.*, the triplets at 7774 and 9263 Å) of the spectrum. Because they are high excitation lines, their use as abundance indicators relies on accurate temperatures and consideration of NLTE effects in abundance analyses. Useful forbidden lines of O I (*e.g.*, those at 6300 and 6363 Å) are very weak ($\sim 5 \text{ mÅ}$ in the Sun) in all but the coolest stars, vanishingly so in metal-poor stars. Lines of OH do exist in the infrared, but their use requires high resolution and S/N observations with Fourier transform spectrographs or newly available infrared echelle spectrographs. Near-UV

lines of OH are also useful for abundance determinations; however, they lie in a crowded region of the stellar spectrum which is near the atmospheric cutoff. CO lines in the optical and infrared can also be used to determine O abundances in cool stars. Careful spectral synthesis is required to infer accurate abundances from often weak and blended lines; additionally, care must be taken in the treatment of molecular equilibria. Because of these difficulties, stellar O abundance determinations are, unfortunately, few in number compared to other metals such as Fe.

3. *This Dissertation*

This dissertation focuses on O abundances as derived mainly from the 7774 Å O I triplet. There are several reasons for this choice. The permitted triplet is very strong in solar-type stars. Even in very metal-poor ($[\text{Fe}/\text{H}] \sim -3$) stars, the triplet is expected to maintain a strength of a few mÅ. Other potential spectral features, namely the 6300 Å [O I] line, are weak in even metal-rich stars. Therefore, internal abundance uncertainties will be larger for this line and observations of this line are quite difficult in the most metal-poor dwarfs. It was thought to be important here to concentrate on dwarfs since their O abundances should not be affected by possible O $\rightarrow$ N processing. Additionally, the 7774 Å O I triplet does not require particularly high resolution to observe. Again, this is not the case for other features such as the 6300 Å [O I] line. The weakness of the forbidden line and its proximity to a Sc feature requires high resolution observations to be useful for abundance analysis. Therefore, use of the forbidden line to derive abundances in faint halo dwarfs or faint open cluster dwarfs is difficult at best with moderate or large aperture telescopes.

The focus on the 7774 Å O I abundances is also due to questions which have arisen concerning their usefulness and accuracy. A review of the literature indicates that the permitted triplet seems to yield larger O abundances than other O features. Whether this is due to NLTE departures, complications arising from atmospheric temperature inhomogeneities, or other effects was not clear at the outset of this study. Therefore, even if the triplet was found not to yield useful abundances, it was still hoped that its study might yield interesting information about stellar atmospheres. The goal, of course, was to hopefully resolve these various discrepancies and exploit the advantages that the 7774 Å triplet offers by deriving O abundances for metal-poor dwarfs as well as open cluster dwarfs and subgiants. As we shall now see, it appears that these goals have been reasonably met.

Chapter 2. A Resolution to the 7774 Å O I Abundance Discrepancy

Abstract

We investigate the discrepancy between [O/Fe] ratios of metal-poor stars derived from the 7774 Å O I triplet and [O/Fe] ratios determined from other oxygen lines. We propose a possible resolution to this discrepancy which also eliminates the correlation of [O/Fe] and T_{eff} found in a recent 7774 Å O I analysis. The equivalent widths of Abia & Rebolo (1989) are found to be systematically too high by 25%. Arguments are presented that current temperature estimates for halo stars are 150 – 200 K too low. Using the guidance of both model atmospheres and other empirical color– T_{eff} relations, we construct new color temperature relations for metal-poor stars. These relations are tied to the temperature scale of Saxner & Hammarback (1985) for metal-rich stars. We use $(b - y)$ and $(V - K)$ indices to redetermine values of T_{eff} for a handful halo stars. $(B - V) - T_{\text{eff}}$ relations which do not take into account the effects of metallicity are found to be inadequate. Revised [O/Fe] ratios are determined using the new temperature scale. The mean abundance ratio of the reanalyzed halo dwarfs is [O/Fe] $\sim +0.52$. There is no trend of [O/Fe] with [Fe/H] or T_{eff} .

1. The 7774 Å O I Discrepancy

The results of Clegg, Lambert & Tomkin (1981) and Andersen *et al.* (1988), who derive O abundances from the 7774 Å O I triplet, show that [O/Fe] $\sim -0.5 \times [\text{Fe}/\text{H}]$ for Population I dwarfs having $-1 \leq [\text{Fe}/\text{H}] \leq 0$. The results of Conti *et al.* (1967) were the first which indicated the existence of supersolar [O/Fe] ratios for metal-poor stars. This was later confirmed by the work of Sneden, Lambert & Whitaker (1979) who derived O abundances in metal-poor dwarfs using the 7774 Å triplet. They found [O/Fe] $\sim +0.5$ in halo field stars, though with appreciable star-to-star scatter. In an important paper, Abia & Rebolo (1989, hereafter AR) determined O abundances for halo dwarfs and subgiants using the 7774 Å O I triplet. Their surprising results indicated that [O/Fe] ~ 0.75 at [Fe/H] ~ -1.0 and *increases* to [O/Fe] $\sim +1.2$ at [Fe/H] ≤ -2.5 . Their data also demonstrate appreciable star-to-star scatter. Recently, Tomkin *et al.* (1992, hereafter TLIS) have derived O abundances in halo dwarfs using the 7774 Å triplet. The mean [O/Fe] ratio in the range $-2.9 \leq [\text{Fe}/\text{H}] \leq -1.1$ is $+0.80$ with significant scatter. Statistical analysis shows that there is no correlation between [O/Fe] and [Fe/H] at greater than the 82% confidence level in their sample.

Gratton & Ortolani (1986) and Barbuy (1988) have determined the O abundance in halo giants using the 6300 Å [O I] line. Their data yield a constant [O/Fe] ratio of approximately $+0.35$ with

little scatter. This value is also in agreement with the $[\text{O}/\text{Fe}]$ ratio derived from the 6300 Å line in globular cluster giants (*e.g.*, Sneden *et al.* 1991). This review of past work indicates that analyses of metal-poor dwarfs using the 7774 Å O I triplet yield significantly larger O abundances than similar analyses of metal-poor giants using the 6300 Å [O I] line. Several possible explanations of this discrepancy have been proposed.

An immediate concern is that the giants' O abundances may have been affected by stellar nucleosynthesis and subsequent mixing. O→N cycling and subsequent dredge up should reduce the surface O abundance of giants. Indeed, some giants are observed to have $[\text{O}/\text{Fe}] < 0$, but also show the expected concomitant N overabundance. Evaluation of the mixing explanation requires 6300 Å [O I] observations of the giants' progenitors— *i.e.*, halo dwarfs. Such observations have been presented recently by Spite & Spite (1991). The mean $[\text{O}/\text{Fe}]$ ratio of three dwarfs with $[\text{Fe}/\text{H}] \leq -1.0$ is +0.36. Spiesman & Wallerstein (1991) have determined the $[\text{O}/\text{Fe}]$ ratio in two metal-poor K dwarfs using the 6300 Å [O I] line. They find $[\text{O}/\text{Fe}]$ values of +0.3 and +0.4. Hence, the discrepancy is not just one between giant and dwarf abundances.

Bessell, Sutherland & Ruan (1991) have derived O abundances in 6 metal-poor dwarfs and 2 giants using OH lines in the near UV (~ 3150 Å). The mean¹ $[\text{O}/\text{Fe}]$ ratio is $\sim +0.5$. The resulting linear correlation coefficient indicates that there is a trend of $[\text{O}/\text{Fe}]$ with $[\text{Fe}/\text{H}]$ at the 96.5% confidence level. In any case, their mean abundance is less than that typically derived for halo stars from the O I 7774 Å triplet. Thus, the discrepancy is not just between the O I triplet and the 6300 Å forbidden line.

The evidence presented in this section suggests that the discrepancy in $[\text{O}/\text{Fe}]$ ratios for metal-poor stars is solely due to abundances derived from the O I 7774 Å triplet. This same conclusion was reached by TLLS, whose data show a definite trend with T_{eff} . They argue that deficiencies in the analysis of the 7774 Å O I triplet may arise from shortcomings in the model atmospheres or in their NLTE analysis. It has been shown that the O I triplet is formed under NLTE conditions in the Sun (Altrock 1968). These NLTE effects are expected to become larger at lower values of $[\text{Fe}/\text{H}]$ and $\log g$ (Kiselman 1991). However, NLTE abundance calculations (*e.g.*, AR, Kiselman 1991, and TLLS) do not seem to significantly affect the O abundances of metal-poor stars derived from the O

¹ We computed the mean using their $[\text{O}/\text{Fe}]$ value for HD 140283 determined with $\log g = 3.3$ as this “low” value is preferred in recent abundance analyses by several authors. Using the O abundance consistency argument presented by Bessell *et al.* (1991) which lead them to favor $\log g = 4.5$ would be begging the question we seek to address in this paper. However, it must be acknowledged that the $\log g$ value for this important star is subject to some uncertainty.

I triplet. I present here a possible resolution to the 7774 Å O I discrepancy and provide evidence to support this resolution.

2. 7774 Å O I Equivalent Widths

We begin by examining the 7774 Å O I triplet equivalent widths themselves. As part of his thesis work, the author has obtained data (some kindly provided by Dr. Ann Boesgaard) for several of the stars observed by AR— though, at the time, no effort was being made to systematically check on AR’s results. Full descriptions of the observations and reductions will be presented in future chapters in this dissertation. Shortly after these first data were analyzed, it was clear that large systematic differences existed with the equivalent width measurements of AR for the few of their stars that were observed. An intentional effort was then made to observe more of their stars. The systematic differences persisted and it was found that other investigators’ data confirmed them.

Table 1 compares equivalent width measurements (all values are in mÅ) of stars in common between AR and other investigators. The first column lists the star. The second column gives the equivalent width of the whole triplet as measured by AR. Lower limits indicate stars in which not all of the lines of the triplet were measured. The corresponding equivalent widths of TLLS are listed in column three.

Column four of Table 1 gives the equivalent width measured by the author from spectra taken with the Faint Object Spectrograph on the University of Hawai’i 2.2-m telescope on Mauna Kea. Again, full details of the observations and reductions will be presented later. Briefly, though, the data were obtained using a 1200 l/mm grating blazed at 6300 Å. The detector was a 1024×1024 Tektronix CCD. The spectra have a measured resolution of ~ 1.2 Å and all have S/N ~ 400 . Dr. A. Boesgaard kindly provided CCD spectra acquired at Palomar of several metal-poor stars. These spectra have a resolution of 0.21 Å and a typical S/N of 200. The measured equivalent widths are listed in column five of Table 1. The sixth column lists values from Sneden *et al.* (1979).

We find excellent agreement between the FOS and Palomar measurements and those given by TLLS and Sneden *et al.* (1979). This comparison uses additional stars in common that are not listed in Table 1 since they were not observed by AR. The mean difference between the AR measurements and the other measurements (where we prefer higher resolution observations in case of overlap), in the the sense

$$[\text{EW}(\text{AR}) - \text{EW}(\text{Other})]/\text{EW}(\text{Other})$$

is found to be +23.9%. There is no correlation of the differences with [Fe/H] and any such trend

can be excluded as the cause of the increase of $[\text{O}/\text{Fe}]$ with decreasing $[\text{Fe}/\text{H}]$ seen in AR’s data. The systematic difference between AR’s equivalent widths and four other sets of data which are consistent with one another strongly suggests that their values are systematically too large. We can offer no explanation as to why their values are apparently in error. Using (what appear to be) more secure 7774 Å O I equivalent widths results in a significant reduction in $[\text{O}/\text{Fe}]$ for most of AR’s stars.

3. The Halo Star Temperature Scale

Because the 7774 Å triplet is composed of high-excitation (9.15 eV) lines, accurate temperatures are needed to derive secure abundances. A systematic underestimate of stellar effective temperatures could produce significant 7774 Å O I overabundances. TLLS did not favor this explanation as a cause of the 7774 Å O I discrepancy because the 0.6 dex difference between the abundances of AR and those derived from, *e.g.*, the 6300 Å [O I] line, would require (typically) the T_{eff} values to be underestimated by ~ 500 K. However, we note that any underestimate need *not* be this extreme to account for the large $[\text{O}/\text{Fe}]$ ratios found by AR. Their systematically larger equivalent widths account for a significant portion of the difference. The adoption of a better determined solar iron abundance, $\log N(\text{Fe})_{\odot} = 7.50$ (a value indicated by several pieces of recent evidence and consistent with the meteoritic abundance), lowers the $[\text{O}/\text{Fe}]$ values for many stars. Any systematic underestimate of T_{eff} may also raise the stellar iron abundance, thus lowering the $[\text{O}/\text{Fe}]$ ratio. For many of AR’s stars, it is in fact rather easy to lower $[\text{O}/\text{Fe}]$ by ~ 0.5 dex if these factors work in tandem. We have already showed that a systematic overestimate of the 7774 Å O I line strength by AR apparently exists. Below, we make our case for the other effects mentioned.

3.1 Color– T_{eff} Relations Guided by Model Atmospheres

Synthetic colors can be determined from model atmospheres by the calculation of flux ratios over various ranges in wavelength. By associating the calculated color with the effective temperature characterizing a specific model, theoretical color– T_{eff} relations can be constructed. Here, we investigate such relations using the new extensive models of Kurucz (1992). A preliminary description of these models is provided by Kurucz (1991). These models are a significant improvement over the Kurucz (1979) models. The number of lines included in the new models stands at 58 million, a factor of 34 more than the 1979 models. Because of this, the “missing opacity” problem has been completely removed (Kurucz 1991). The new models have more layers than the old models and

reach deeper in the stellar atmosphere. The new models also have more wavelength points for flux computation and use a more sophisticated treatment of convection.

We adopt the premise that theoretical color– T_{eff} relations may be incorrect in an *absolute* sense— *i.e.*, that the zero-point calibrations of various theoretical colors may be in error. However, we assume that the *relative* colors and temperatures (*i.e.*, having differences due to $\log g$, $[\text{Fe}/\text{H}]$, *etc.*) derived from the models are correct. In order to fix the color– T_{eff} relations for metal-poor stars in an absolute sense we use the semi-empirical guidance of the Saxner & Hammarback (1985, hereafter SH) relations. These relations were chosen because the scatter is relatively small (± 60 K) as is their estimate of possible systematic error (± 75 K). The SH calibrations are based on the infrared flux method which uses estimated angular diameters and observed integrated fluxes. Because the Planck function is only weakly dependent on temperature in the infrared, the SH results are virtually free of model dependences. SH also provide improved zero-points for the magnitudes in the Johnson JKL photometric system. We consider various temperature sensitive color indices, in turn, below.

3.1.1 Strömgren $(b - y)$

The semi-empirical $(b - y)$ – T_{eff} relation found by SH is given by

$$T_{\text{eff}} = 8290 - 6290(b - y)\{1 - 0.108[\text{Fe}/\text{H}]\}$$

which is valid for $0.20 \leq (b - y) \leq 0.40$ and $-0.50 \leq [\text{Fe}/\text{H}] \leq +0.20$. They found no dependence of T_{eff} on $\log g$. This is borne out by comparison with the Kurucz model relations, which (for $[\text{Fe}/\text{H}] \leq 0$) anyway, show minimal dependence (± 30 K) on $\log g$ over the temperature range of interest here ($\sim 5000 - 6400$ K). In Figure 1, the $(b - y)$ vs. T_{eff} plane is plotted for Kurucz models having a range of metallicities ($\log g = 4.0$ for all the models plotted). It is apparent that, at $[\text{Fe}/\text{H}] \sim -0.5$, the relations essentially run together for our T_{eff} range of interest. Even at 5000 K, the possible error in assuming a single relation for metal-poor stars is only 40 – 50 K. It seems, then, that the SH $(b - y)$ – T_{eff} relation for $[\text{Fe}/\text{H}] = -0.5$ (the lower limit at which the relation was claimed to be valid by them) can be used for all metal-poor stars (having the same reasonable range in $\log g$ of course).

Figure 2 shows the $(b - y)$ relation of SH compared to that of Magain (1987a, hereafter M87), which is commonly used in abundance analyses of metal-poor stars. Magain uses the same approach and calibrations of SH, essentially extending their work to Pop II stars. However, as noted by M87, the final T_{eff} values are not on the same scale. Figure 2 indicates that there is a roughly constant

difference of ~ 145 K between the temperatures derived from the relation of M87 and SH. This is in the sense that the M87 relation, which is assumed to be unique for $[\text{Fe}/\text{H}] \leq -1.0$, yields lower values of T_{eff} .

3.1.2 Johnson ($B - V$)

The semi-empirical $(B - V) - T_{\text{eff}}$ relation found by SH is given by

$$T_{\text{eff}} = 8065 - 3580(B - V)\{1 - 0.196[\text{Fe}/\text{H}]\}$$

which is valid for $0.30 \leq (B - V) \leq 0.63$ and $-0.50 \leq [\text{Fe}/\text{H}] \leq +0.20$. The synthetic colors from the Kurucz models indicate that the $(B - V) - T_{\text{eff}}$ relation shows negligible sensitivity to $\log g$ for $[\text{Fe}/\text{H}] \gtrsim -0.5$. For lower metallicities, however, the effects of $\log g$ are seen to become more important. At $[\text{Fe}/\text{H}] = -2.5$ and $(B - V) = 0.40$, the difference between $T(\log g = 3.5)$ and $T(\log g = 4.5)$ is ~ 85 K. For metal-poor stars it may be beneficial to take this $\log g$ dependency into account using the predictions of the Kurucz models.

Figure 3 shows the metallicity sensitivity of the Kurucz $(B - V) - T_{\text{eff}}$ relations. Even for $[\text{Fe}/\text{H}] \leq -2.0$ *there is still sensitivity to the metal abundance*. This finding questions the validity of similar relations, such as that of M87, which include no metallicity dependence. In fact, this metallicity dependence may be visible in M87's figure 1, where the lowest metallicity point does lie significantly below the mean trend.

We construct an improved $(B - V) - T_{\text{eff}}$ relation for metal-poor stars by forcing the Kurucz $[\text{Fe}/\text{H}] = -0.5$ model relation to match the SH relation for $[\text{Fe}/\text{H}] = -0.5$ at $(B - V) \sim +0.6$. The $(B - V) - T_{\text{eff}}$ relation for an arbitrary metallicity, $[\text{Fe}/\text{H}] \leq -0.5$, is then determined from the *relative* difference between the Kurucz model relations (at $[\text{Fe}/\text{H}] = -0.5$ and at the $[\text{Fe}/\text{H}]$ in question) and the *modified* Kurucz relation for $[\text{Fe}/\text{H}] = -0.5$ which has been forced to roughly coincide with that of SH. Figure 4 shows these $(B - V) - T_{\text{eff}}$ relations for $[\text{Fe}/\text{H}] = -1.0$ and $[\text{Fe}/\text{H}] = -2.5$ compared to the relation of M87. Temperatures for higher metallicity halo stars would be underestimated (in some cases by ≥ 200 K) by the Magain relation compared to the method devised here.

3.1.3 Johnson ($V - K$)

$(V - K) - T_{\text{eff}}$ relations are often used to determine stellar temperatures for abundance analyses. Relations used include those of M87 and Carney (1983, hereafter C83), both of which contain no dependence on metallicity. The Carney relations are based on optical and near infrared spectrophotometry of halo dwarfs. The spectrophotometry is converted to surface flux using the Hayes &

Latham (1975) Vega calibration. The surface fluxes are compared to those from older ATLAS6 models in order to derive temperatures. Figure 5 shows the metallicity dependence of the $(V - K) - T_{\text{eff}}$ relations given by the Kurucz (1992) models. While the dependence on metallicity is small, it is not negligible (according to the Kurucz models, the dependence on $\log g$ is negligible for $V - K$ at all metallicities and over the T_{eff} range of interest). By using the relations of C83 or M87, the Kurucz models indicate that the temperature would be increasingly underestimated with decreasing $[\text{Fe}/\text{H}]$. It is interesting that the $[\text{O}/\text{Fe}]$ ratios of AR, who used (in part) the $(V - K)$ relations of C83, increase with decreasing metallicity in the halo. Such behavior could be introduced as part of the abundance analysis if the temperature values were systematically underestimated with decreasing $[\text{Fe}/\text{H}]$.

3.2 Consistency of Color-Temperature Relations

If the behavior of color- T_{eff} relations predicted by the Kurucz models is correct, then we would expect that the difference in T_{eff} derived from $(b - y)$ and $(B - V)$ (as well as the difference between T_{eff} derived from $(b - y)$ and $(V - K)$) using relations such as those given by M87 should be a function of metallicity. Table 2 gives photometry for metal-poor dwarfs (or subgiants) included in the C and O abundance study of TLLS. We use only those stars which have photometry available in C83 (as this is a primary source of $(V - K)$ data for halo dwarfs). Only stars whose colors are within the range covered by the M87 relations are used. A few additional stars not analyzed by TLLS are included to round out the $[\text{Fe}/\text{H}]$ coverage and increase the sample size.

Photometry was taken from C83 or from Laird, Carney & Latham (1988) if more precise values were available there. Some of the $(b - y)$ values of C83 are actually $(B - Y)$ values given by Eggen (1978). For these stars, we used the $(b - y)$ photometry from Hauck & Mermilliod (1990). For BD+21 607 no $b - y$ value was available so we retain the $(B - Y)$ value. Reddening corrections were applied based on the $E(B - V)$ values listed in Laird *et al.* (1988) and Ryan & Norris (1991b). The relations $E(b - y) = 0.7E(B - V)$ and $E(V - K) = E(B - V)/0.36$ were assumed. Most metallicities were taken from TLLS; for those stars not studied by TLLS, values were adopted based on literature estimates.

Using the relations of M87, effective temperatures based on $(b - y)$, $(B - V)$, and $(V - K)$ were derived. The difference $T(b - y) - T(B - V)$ *vs.* $[\text{Fe}/\text{H}]$ is shown in Figure 6. We find the Spearman coefficient for $[\text{Fe}/\text{H}]$ and $\delta T = T(b - y) - T(B - V)$ to be $r_s = 0.755$. Hence the correlation seen in Figure 6 is significant at $> 99.9\%$ confidence level and is in the sense that the Kurucz models predict.

Clearly, $(B - V) - T_{\text{eff}}$ relations which do not take metallicity and surface gravity into account can lead to large, systematic errors.

Figure 7a is the analogous plot with the temperature difference $T(b - y) - T(V - K)$. There is a suggestion of a correlation in this figure and in the opposite sense from Figure 6 as predicted by the Kurucz models. The ordinary correlation coefficient is $r = -0.350$ which is, at best, only marginally significant at the 91% confidence level. Indeed, $r_s = -0.266$ which is significant at only $\sim 80\%$ confidence level. To investigate the matter further, we added additional points at higher metallicities to increase the $[\text{Fe}/\text{H}]$ baseline. We argue this is a valid approach because: a) the Kurucz models indicate that the $(b - y) - T_{\text{eff}}$ relations are not affected by metallicity for $[\text{Fe}/\text{H}]$ as high as -0.5 b) Although M87 restricts the use of his relation to $[\text{Fe}/\text{H}] \leq -1.0$, it is essentially equivalent to that of C83 for $5100 \leq T_{\text{eff}} \leq 6200$ K. This latter relation has been used by AR in the derivation of O abundances. Carney argues that this relation is independent of metallicity (see his figures 4 and 6). Use of the similar M87 relation is then adequate to test this claim.

We were able to find 5 stars with metallicity determinations in Laird *et al.* (1988) for which $-1.0 < [\text{Fe}/\text{H}] < -0.5$ with photometry in C83. These are also listed in Table 2. Figure 7b shows the temperature difference, $T(b - y) - T(V - K)$ *vs.* $[\text{Fe}/\text{H}]$ when one includes these higher metallicity stars. A real trend with metallicity, again in the direction predicted from the Kurucz model results, is more evident than in Figure 7a. The Spearman coefficient is $r_s = -0.553$, indicating a correlation at $\geq 99.8\%$ confidence level. The temperature difference for HD184499 is particularly large². If one excludes this point as an outlier then $r_s = -0.512$, which still indicates a real correlation at $\geq 99.5\%$ confidence level.

3.3 Spectroscopic Temperature Determinations

3.3.1 T_{eff} from the $I(\lambda 6748)/I(\lambda 6750)$ and $I(\lambda 6748)/I(\lambda 6752)$ Ratios

Boesgaard & Tripicco (1987) found a relation between the line ratios of the $\lambda 6748.87$ Cr I and the $\lambda 6750.16$ Å Fe I lines as well as the $\lambda 6748.87$ and $\lambda 6752.72$ Fe I lines. We have recalibrated their

² Carney (1992, private communication) has confirmed the discrepancy of $T(V - K)$ with respect to other photometric values. He has suggested two explanations for this anomaly: a) a slim possibility that the star HD 184449 (as opposed to HD 184499) was actually observed. HD 184449 is a faint A0 star near the Galactic plane and may be heavily reddened, resulting in a large $(V - K)$ index accidentally attributed to HD 184499 b) HD 184499 could have a red companion which would result in an anomalous $(V - K)$ index.

relation by using temperatures on the SH scale. The rederived relations are given by:

$$T_{\text{eff}} = 1130.23\{I(6748)/I(6750)\} + 5774.37$$

$$T_{\text{eff}} = 535.40\{I(6748)/I(6752)\} + 5712.59$$

The rms residuals of these relations are ± 14.5 and ± 19.2 K.

There is some concern that the application of these equations, involving the line ratio of the Cr I and Fe I lines, to halo stars may be subject to abundance effects. As reviewed in Wheeler *et al.* (1989), all studies of halo field stars have found that $[\text{Cr}/\text{Fe}] \sim 0$ (or perhaps slightly lower), regardless of $[\text{Fe}/\text{H}]$. This result is confirmed by the more recent work of Gratton & Sneden (1991) who find $[\text{Cr}/\text{Fe}] = -0.04 \pm 0.05$ for their sample of metal-poor field stars. This result is particularly robust in terms of the present paper in that the $[\text{Cr}/\text{Fe}]$ ratio is quite insensitive to the adopted T_{eff} value (see Table 4 of Gratton & Sneden 1991). Table 1 of Wheeler *et al.* (1989) indicates that $[\text{Cr}/\text{Fe}] < 0$ in the most metal poor globular clusters ($[\text{Cr}/\text{Fe}] \sim -0.1$ is the average ratio for M15 and M92 as determined by Cohen 1979). Again, this result is seemingly confirmed by the recent work of Gratton & Sneden who find $[\text{Cr}/\text{Fe}] = -0.14 \pm 0.09$ for their sample of 30 giants in 13 globular clusters. However, it is interesting to note that for the cluster stars, Table 4 in Gratton & Sneden (1991) indicates there is significant sensitivity of $[\text{Cr}/\text{Fe}]$ to T_{eff} . The T_{eff} values adopted by them for the cluster giants can be shown to be a bit low in comparison with the present work (*e.g.*, as they directly state, they prefer T_{eff} values some 70 K lower than the values given by the Cohen, Frogel & Persson (1978) relation which we prefer here— see §3.3.5). If one raises their temperatures to account for this difference, the cluster $[\text{Cr}/\text{Fe}]$ ratio is much closer to that found for their field stars. Whether differences in the abundance patterns of globular clusters and the halo field are genuine or due to inconsistent T_{eff} scales will require evaluations, which are not the immediate concern of this paper, on a case-by-case basis. In any case, even if $[\text{Cr}/\text{Fe}]$ is somewhat less than 0 in the halo field, this abundance effect would cause the temperature derived from the relations above to be *underestimated*.

Dr. M. Spite kindly provided four high-resolution spectra of three halo stars: BD +183423, HD 189558, and HD 205650. These were obtained with the coude spectrograph of the CFHT using a reticon detector. The temperatures derived from the measured line ratios are listed in Table 3 along with the temperatures from the M87 relations and the C83 ($V - K$) relation. The spectroscopic temperatures are substantially larger than the photometric estimates.

3.3.2 T_{eff} from the $I(\lambda 6297)/I(\lambda 6290)$ Ratio

As part of another program, observations of the 6300 Å region have been obtained for numerous F and G disk dwarfs and several halo dwarfs. Full details of the observations and reductions will be presented later. Briefly, the observations were obtained with the coudé spectrograph of the UH 2.2-m telescope on Mauna Kea. The detector was a 1024×1024 Tektronix CCD. The measured resolution of the spectra is ~ 0.21 Å. The S/N $\gtrsim 500$ for the disk dwarfs and ~ 250 for the halo dwarfs. Though the 6300 Å region of the stellar spectrum is rich, we could only find 2 lines with suitable characteristics to be potentially good temperature indicators. These are the 6290.97 and 6297.80 Å Fe I lines. The solar equivalent widths are 66 mÅ and 65 mÅ (Moore, Minnaert & Houtgast 1966). These lines were chosen because they are strong enough to be observed in metal-poor stars yet not so strong so as to be dominated by desaturation phenomena. Because they are lines of neutral species, they should not be sensitive to $\log g$. Because both lines are of the same species, there should be no dependence of the derived T_{eff} on abundance.

The line depth ratios, $I(\lambda 6297)/I(\lambda 6290)$, were measured for twelve disk dwarfs with $5700 < T_{\text{eff}} < 6540$. Temperatures on the SH scale were determined (most of the stars are from Balachandran (1990), whose temperatures are on the SH scale). The line ratios were found to be correlated with T_{eff} at $\geq 99.9\%$ confidence level. The two quantities are related by:

$$T_{\text{eff}} = -2438.9\{I(6297)/I(6290)\} + 8643.2.$$

Unfortunately, the scatter is more than one would hope for— the rms deviation between the the fitted and actual temperatures is ± 40 K. The line ratios of two halo stars were then measured and the T_{eff} calculated from the Pop I calibration. We find $T_{\text{eff}} = 5926$ for HD 184499 and $T_{\text{eff}} = 6141$ for HD 201891. These values are listed in Table 3 along with temperatures derived from the relations of Magain and the $(V - K)$ relation of C83. The spectroscopic temperature is much larger than photometric estimates for HD 201891. In the case of HD 184499, the spectroscopic temperature is larger than the estimate from $(b - y)$, but comparable to the estimate from $(V - K)$. This curious discrepancy between the photometric estimates was noted before in §3.2.

3.3.3 Spectroscopic T_{eff} Values in the Literature

Perrin *et al.* (1977) have determined effective temperatures for metal-poor dwarfs and subgiants from spectroscopic analyses. These temperatures are listed in Table 4 along with temperatures derived from the relations of M87 and C83 for stars with $-0.90 \leq [\text{Fe}/\text{H}] \leq -2.00$ according to Perrin *et al.* (1977) (it would seem, however, that several of these stars have $[\text{Fe}/\text{H}] < -2.00$ as determined

from more recent analyses). Comparison of the T_{eff} values in the table shows that the spectroscopic estimates are generally larger than the those derived from existing photometric relations.

3.3.4 A Specific Example: HD 103095

Smith, Lambert & Ruck (1992) have recently conducted a fine analysis of the halo subdwarf HD 103095 (=Groombridge 1830). They employed a self-consistent, iterative analysis of 10 Ca lines and 18 Fe lines to determine that $T_{\text{eff}} = 5170 \pm 70$ K. This result is in good agreement with the spectroscopic temperature, 5140 K, determined by Perrin *et al.* (1977). The value of Smith *et al.* (1992) is substantially larger than the photometric value listed in Table 4 and values (~ 5000 K) typically used in abundance analyses. Correcting the O data of TLLS for a temperature of 5170 K (by use of their Table 5) and adopting $[\text{Fe}/\text{H}] = -1.30$ from the fine analysis of Smith *et al.* (1992) yields $[\text{O}/\text{Fe}] = +0.43$. This is a 0.46 dex reduction from the value found by TLLS.

3.3.5 Spectroscopic Constraints on T_{eff} Estimates

In spectroscopic abundance determinations, one way which errors in the adopted temperature scale can manifest themselves is by a trend between the derived abundance and the excitation potential of the lines. In a recent careful analysis of metal abundances in red giants in the globular cluster M5, Sneden *et al.* (1992) find no such trend in their data (see their figures 3 and 4). Additionally, the star-to-star scatter in the derived $[\text{Fe}/\text{H}]$ values is very small (~ 0.05 dex). These results strongly suggest that the temperature scale they adopt is reasonably correct.

The temperature scale used by Sneden *et al.* (1992) is that of the Cohen, Frogel & Persson (1978, hereafter CFP) $(V - K) - T_{\text{eff}}$ relation. Using the $(V - K)$ values for HD 103095 and HD 87140 given by C83 and the CFP relation (as given in their table 3 and figure 4), temperatures of 5125 and 5148 K, respectively, are found. These values are substantially larger than those adopted by TLLS, but in good agreement with the temperatures rederived in §4.1. As discussed in §4, the higher temperatures we find, supported in part by the findings of Sneden *et al.* (1992), reduce the derived $[\text{O}/\text{Fe}]$ ratio significantly and support the notion that temperature scales commonly adopted for metal-poor dwarfs may be too low.

Luck & Bond (1981, 1985) have determined spectroscopic temperatures for a good number of red giants by forcing different Fe I lines of different excitation potential to yield the same abundance. In Table 5, we list their stars also having $(V - K)$ photometry in C83. The second column gives Luck & Bond's spectroscopic temperature estimate while the third column lists the photometric temperature

that would be found on the scale of CFP. In the third column are photometric temperatures that would be found using the relation of C83. For four of the six stars, the spectroscopic temperatures are substantially larger than the photometric estimates using the C83 relation, but are in good agreement with temperatures of CFP. Moreover, the T_{eff} estimates for all the stars (no T_{eff} estimate on the CFP scale could be made for HD 184266 since it lies outside their calibration) on the CFP scale are systematically ~ 160 K higher than the values on the C83 scale. The results of Luck & Bond (1981, 1985) and Sneden *et al.* (1992) strongly suggest that the temperature scale of CFP is correct. The scale of the C83 relation is essentially equivalent to that of M87 and both relations have been used in the determination of O I (and other elements') abundances from the 7774 Å triplet. The large discrepancy between the CFP and C83/M87 scales again suggests that previous analyses of metal-poor stars have used values of T_{eff} that are too low and have thus overestimated the true [O/Fe] ratio.

One must, however, be aware of a very important caveat concerning the above arguments about the constraints provided by spectroscopy on stellar T_{eff} estimates. The spectroscopy noted in this section is that performed on *red giants*. The atmospheres of intrinsically luminous giants near the tip of the RGB in globular clusters and the atmosphere of an intrinsically faint dwarf such as HD 103095 (= Gmb 1830) are vastly different. It must be asked whether the CFP ($V - K$) T_{eff} scale for giants can be applied to dwarfs. An answer to this important question can be obtained using the Kurucz (1992) models. They indicate that the ($V - K$) – T_{eff} relation is insensitive to $\log g$. For example, at the ($V - K$) color of HD 103095, models having $\log g$ of 1.5 and 4.5 dex differ in T_{eff} by ≤ 10 K.

4. A Reanalysis of 7774 Å O Abundances

To investigate the effect of a different temperature scale on O abundances derived from the 7774 Å line, we reanalyzed the data of TLLS. We selected 14 of their stars to provide a range in temperature and metallicity. We also included HD 201891, using the Palomar and FOS equivalent width measurements given in Table 1. The stars are listed in the first column of Table 6.

4.1 New Temperatures

Temperatures for stars having $0.30 \lesssim (b - y) \lesssim 0.43$ were determined using the $(b - y)$ index. Since the Kurucz models show that the $(b - y)$ – T_{eff} relation for $[\text{Fe}/\text{H}] \leq -0.5$ is independent of $[\text{Fe}/\text{H}]$, we simply adjust the zero-point of the Magain relation to bring it into agreement with the SH relation

shown in Figure 2. The resulting relation is given by:

$$T_{\text{eff}} = 8335 - 6670(b - y).$$

For cooler stars (*i.e.*, those with $(b - y) \gtrsim 0.43$), T_{eff} was determined using the $(V - K)$ index. This was done for two reasons: a) the empirical calibration of SH does not extend to cool stars and b) the Kurucz models show increased sensitivity of the $(b - y) - T_{\text{eff}}$ relation to $[\text{Fe}/\text{H}]$ in cool stars. We first determined temperatures on the SH scale for the Hyades stars given in C83. This was done using the $\text{H}\beta$ and $(b - y)$ values from Hauck & Mermilliod (1990) and the $(B - V)$ data of C83. The resulting T_{eff} values are plotted versus C83's $(V - K)$ data in Figure 8. The solid line is the fit to the data given by:

$$T_{\text{eff}} = 8145.5 - 1539.8(V - K).$$

The rms scatter is ~ 12 K.

This relation has been calibrated using only the data with $0.95 \leq (V - K) \leq 1.50$ because of the limited range where the SH calibration is valid. To use the $(V - K)$ index as a temperature indicator for cooler stars, we simply mate the Carney relation to the one above at $(V - K) = 1.50$. Since this is essentially just a zero-point shift one could, in principle, mate the Kurucz model relations to the relation above at $(V - K) = 1.50$. However, we prefer to use the Carney relation because: a) it is desirable to use relations with some empirical guidance and b) mating the Kurucz relations would yield even larger temperatures than those found by mating the Carney relation— in the interest of being conservative, we choose the Carney relation. As a final step, we add the temperature difference due to the metallicity difference between the adopted relation (assumed to be valid for a Hyades metallicity which we take here as $[\text{Fe}/\text{H}] = +0.10$) and the star in question. This is computed via the use of the Kurucz models. This final step is not crucial as it only affects the derived T_{eff} by ≤ 30 K or so due to the small metallicity sensitivity for $(V - K) \geq 1.50$.

To test the consistency of the temperatures derived here we compared the temperatures derived from the $(b - y)$ index with those derived from the $(V - K)$ index for those stars with $(b - y) \lesssim 0.43$ and $(V - K) \gtrsim 1.50$. For the five such stars listed in Table 6, the mean temperature difference, $T(b - y) - T(V - K)$, is -14.6 K with a 1σ dispersion of ± 64 K. This good agreement, well within the uncertainties, lends some credence to our revised temperatures. For the stars in Table 6 with bluer colors, the temperature difference becomes substantially larger, -65 K with similar dispersion. While this difference and dispersion is not larger than the uncertainties, it could be explained if the metallicity sensitivity of the $(V - K) - T_{\text{eff}}$ relation as predicted by the Kurucz models is overestimated. The temperature difference for the bluer stars in Table 6 calculated without adjusting the $(V - K)$ temperature for the metallicity sensitivity using the Kurucz models is found to

be $+35 \pm 19$ K— an improvement. This raises the question of the reality of the metallicity sensitivity of $(V - K) - T_{\text{eff}}$ relations. Though the results of §3.2 indicated that there does appear to be such sensitivity at a level consistent with the Kurucz models’ predictions, the question should be regarded as still open. We stress, however, that our final approach avoids these uncertainties to a large extent by using the $(b - y)$ or the $(V - K)$ index in color regimes where there is no expected metallicity sensitivity. In the region of overlap there appears to be excellent agreement between the temperatures derived from both indices. The final temperatures are shown in column two of Table 6. For HD 103095, we use the temperature from the fine analysis of Smith *et al.* (1992), $T_{\text{eff}} = 5170$. The temperature that would have been derived from our $(V - K)$ relation is 5150, in excellent agreement with the spectroscopic temperature. The temperatures derived here are typically ~ 170 K hotter than the temperatures adopted by TLLS. The difference is larger for the cooler stars.

4.2 New $[\text{Fe}/\text{H}]$ values

Revised temperatures also necessitate the estimation of new $[\text{Fe}/\text{H}]$ values. This was done by first searching the literature for $[\text{Fe}/\text{H}]$ determinations of the stars in Table 6. We preferred those values derived from high resolution spectra taken with CCD detectors. Corrections were first made to bring the $[\text{Fe}/\text{H}]$ values onto a common scale. This was done by adopting a solar iron abundance of $\log N(\text{Fe})_{\odot} = 7.50$ (on a scale with $\log N(\text{H}) = 12$) as indicated by several recent pieces of work (Biemont *et al.* 1991; Holweger *et al.* 1991; Holweger, Heise & Kock 1990). The new values of $[\text{Fe}/\text{H}]$ were then revised according to the temperature sensitivity listed in the specific references. Several sources claim to provide precise abundances, but did not make their abundance scale or temperature sensitivity specifically known. Values from these sources were, in all but a few cases, discarded. The final adopted $[\text{Fe}/\text{H}]$ values are listed in column 3 of Table 6. The fourth column gives the references used in estimating $[\text{Fe}/\text{H}]$. The resulting values are higher than those of TLLS. This is especially so for the coolest stars because: a) the temperatures for these stars are significantly larger than those of TLLS and b) $[\text{Fe}/\text{H}]$ values for the cooler stars taken from other sources seem to be larger than those of TLLS.

4.3 New $[\text{O}/\text{Fe}]$ values

Using the new temperatures, we calculated the $[\text{O}/\text{H}]$ ratios of the stars listed in Table 6. This was done by adjusting the LTE $[\text{O}/\text{H}]$ ratio of TLLS by an amount corresponding to the differences between the T_{eff} values given in Table 6 and those originally used by TLLS. This differential abun-

dance correction was calculated using Dr. M. Spite’s LTE analysis software RAI10. The corrections are in excellent agreement with those listed in Table 5 of TLLS. For HD 201891, which was not studied by TLLS, $\log g = 4.4$ was assumed. The revised LTE [O/Fe] ratio is listed in column 5 of Table 6. Column 6 lists the LTE abundance in column 5 minus the difference between the LTE and NLTE abundances determined by TLLS. We denote this as the NLTE abundance, but this is not totally correct. Because we have adopted different temperatures, the formal NLTE abundance corrections may be different than those found by TLLS. However, our values do indicate, to first order, the magnitude and direction of NLTE effects on the derived abundances.

5. Discussion

Figure 9 shows [O/Fe] *vs.* [Fe/H] using the rederived LTE abundances in Table 5. The mean abundance is [O/Fe] = +0.53 with a 1σ dispersion of ± 0.12 dex. If one compares our Figure 9 to Figure 2 in Bessell *et al.* (1991), it is evident that there is no difference, within the uncertainties, between abundances derived there from the forbidden [O I] line at 6300 Å and the abundances rederived here from the 7774 Å triplet. There would be some revision (~ -0.05 dex) of the [O/Fe] ratio determined from the 6300 Å line due to the temperature scale used here. However, the [Fe/H] values used in the halo data shown in Figure 2 of Bessell *et al.* (1991) are either based on sources using photometric calibrations or on sources which do not make the temperature sensitivity of the [Fe/H] determinations explicitly known. Therefore, reanalysis of this data is difficult; fortunately, the effects of a modified temperature scale on these [O/Fe] ratios are small.

No trend of [O/Fe] with [Fe/H] is visible in the (halo star) data of Figure 9. This is confirmed by statistical analysis. The lack of a trend in [O/Fe] with [Fe/H] conflicts with the findings of AR (1989). Since their equivalent width data seem to be in error, we place more confidence in the results found here. The lack of a trend does not seem to be the result of adopting a new temperature scale. The original results of TLLS also show no trend of [O/Fe] with [Fe/H] in the halo. Our results are consistent with those using [O I] abundances. While the best evidence seems to be that there is no trend of [O/Fe] with [Fe/H] in halo stars observed so far, such trends in the extreme halo, predicted by some chemical evolution models, remain an open question.

Figure 10a shows the LTE [O/Fe] ratios of TLLS versus their T_{eff} values. Also included is HD 201891 for which we calculated an abundance using a temperature ($T_{\text{eff}} \sim 5850$ K) roughly on the temperature scale of TLLS. This subsample has a Spearman coefficient of $r_S = -0.796$, indicating a correlation at $> 99.9\%$ confidence level (The simple linear correlation coefficient indicates a

correlation at the same high confidence level). TLLS noted that such a correlation between their derived $[\text{O}/\text{Fe}]$ ratios and temperatures was present in their sample. We confirm that this correlation persists in the subsample of their stars selected here. Figure 10b shows the LTE $[\text{O}/\text{Fe}]$ ratios determined here versus the revised temperatures. There is no apparent trend of the $[\text{O}/\text{Fe}]$ ratio with T_{eff} in this figure. The value of the Spearman coefficient, $r_s = -0.418$, indicates a trend at only the 87.5% confidence level (the simple linear correlation coefficient indicates a correlation at only the 67% confidence level).

6. Summary and Conclusions

This paper presented a possible resolution to the discrepancy between $[\text{O}/\text{Fe}]$ ratios in halo dwarfs as determined from the 7774 Å O I triplet and the $[\text{O}/\text{Fe}]$ ratios in giants and dwarfs determined from other O features. We find that AR's 7774 Å O I triplet equivalent width measurements are systematically too large by $\sim 25\%$. The smaller 7774 Å O I equivalent widths determined by three different observers using three different observational facilities strongly suggest that it is the AR measurements which are in error.

As has been noted by other workers, the high-excitation 7774 Å lines require accurate stellar temperatures to yield secure abundances. We propose that the temperatures adopted by AR and TLLS in their analyses of the 7774 Å triplet in metal poor stars are 150 – 200 K too low. If this is the case, then the derived O abundances would be overestimated while the derived Fe abundances would be underestimated— leading to artificially high $[\text{O}/\text{Fe}]$ ratios.

We use the theoretical guidance of the extensive new Kurucz (1992) model atmospheres and the SH color– T_{eff} relations for metal-rich stars to construct new, self-consistent color– T_{eff} relations for metal-poor stars. These are compared to the corresponding relations of M87 and C83, which are commonly used to derive temperatures for spectroscopic abundance analyses (including the work of AR and TLLS). We find the following:

- (a) Based on the behavior, as predicted by the Kurucz models, of the SH $(b - y) - T_{\text{eff}}$ relations at low metallicity, the $(b - y) - T_{\text{eff}}$ relation of M87 is found to yield temperatures ~ 150 K too low.
- (b) The Kurucz models predict that $(B - V) - T_{\text{eff}}$ relations are still subject to substantial metallicity dependence— even for $[\text{Fe}/\text{H}] \leq -2$. This prediction is supported by the observed correlation between $T(b - y) - T(B - V)$, as determined from the relations of M87, and $[\text{Fe}/\text{H}]$. $(B - V)$ should be avoided as a temperature indicator for metal-poor stars in the temperature range of interest here (5000-6400 K), unless correction for the metallicity dependence (perhaps using the Kurucz models

in a relative sense) can be made.

(c) The Kurucz models predict a small, but non-negligible, sensitivity of $(V - K) - T_{\text{eff}}$ relations to $[\text{Fe}/\text{H}]$. The correlation of $T(b - y) - T(V - K)$, again determined from the M87 relations, with $[\text{Fe}/\text{H}]$ over two decades of metallicity is consistent with this prediction. This finding seems to be at odds with the current belief (*e.g.*, C83) that there is no such metallicity sensitivity. We know of no reason why the sample of stars (being mainly drawn from TLLS and supplemented with a few stars in our temperature range of interest in order to expand the $[\text{Fe}/\text{H}]$ coverage) in Figure 7b would be biased to show $[\text{Fe}/\text{H}]$ sensitivity. Still, the results of §4.1 indicated that the $[\text{Fe}/\text{H}]$ sensitivity of the Kurucz model $(V - K) - T_{\text{eff}}$ relations may be overestimated. This question remains open and will require additional work.

Spectroscopic temperatures derived from the line ratios $I(\lambda 6748)/I(\lambda 6750)$, $I(\lambda 6748)/I(\lambda 6752)$, and $I(\lambda 6297)/I(\lambda 6290)$ are consistently larger than photometric temperatures calculated from the commonly used relations of M87 and C83. We caution that our spectroscopic temperatures are uncertain. The spectra used to measure the first two line ratios have lower S/N and resolution than is desirable for the purpose at hand. The temperature calibration involving the last line ratio shows appreciable scatter. The important point is that our spectroscopic values are consistently higher than photometric values—this result cannot be due to random errors.

Spectroscopic values of T_{eff} taken from the literature (*e.g.*, Perrin *et al.*, 1977 and Luck & Bond, 1981 & 1985) also appear to be consistently larger than corresponding photometric values. The recent fine spectroscopic analysis of the important subdwarf HD 103095 by Smith *et al.* (1992) yields a temperature 170 K larger than the photometric estimate used by TLLS to determine the O I abundance from the 7774 Å triplet. The spectroscopic analyses of metal abundances in globular cluster red giants by Sneden *et al.* (1992) suggest that the temperature scale of CFP is correct. This scale is ~ 160 K hotter than the M87 and C83 scales, both of which are essentially identical to the temperature scale of AR and TLLS.

We chose a subsample of metal-poor stars from TLLS to study the effects of revised T_{eff} values on the derived $[\text{O}/\text{Fe}]$ ratio. We use a modified version of the M87 $(b - y) - T_{\text{eff}}$ relation on the SH temperature scale to determine T_{eff} for stars having $0.30 \lesssim (b - y) \lesssim 0.43$. We placed the C83 empirical $(V - K) - T_{\text{eff}}$ relation on the SH scale by applying a zero-point shift. This shift was determined from photometric temperatures of Hyades stars on the SH scale. The modified relation was used to determine T_{eff} for stars with $(b - y) \gtrsim 0.43$. These color regimes were selected so that the metallicity sensitivity of the modified color- T_{eff} relations is negligible according to the new Kurucz (1992) models. In a small region of overlap, the two relations yield temperatures in excellent agreement.

The estimated 1σ uncertainty is 65 K, essentially that of the original SH relations for metal-rich stars. The resulting O and Fe abundances (where the Fe values are all on a consistent scale with $\log N(\text{Fe})_{\odot} = 7.50$) yield a mean $[\text{O}/\text{Fe}]$ ratio of +0.53 with a 1σ dispersion of 0.12 dex. This value does not differ, within the respective uncertainties, from the ratios derived from the 6300 Å [O I] line in halo field giants. The equality of the $[\text{O}/\text{Fe}]$ ratios derived from the 7774 Å and 6300 Å lines seems to imply that NLTE corrections computed by AR and TLLS are correct and that inclusion of such effects do not significantly alter LTE O abundances derived from the 7774 Å lines in metal-poor dwarfs in the temperature range considered here. In contrast to the surprising results of AR, there is no trend of $[\text{O}/\text{Fe}]$ with $[\text{Fe}/\text{H}]$. Additionally, no trend of $[\text{O}/\text{Fe}]$ with T_{eff} is seen, thus removing the disturbing such correlation noted by TLLS in their data.

Bessell *et al.* (1991) suggested that deficiencies in the deeper layers of existing model atmospheres may be the cause of the 7774 O I discrepancy. It is very important to note that if this is the case, then altering the effective temperature scale of metal-poor stars may remove the discrepancy but not be the true solution *per se*. The effect of altering the temperature scale would be that abundances derived from other lines formed higher up in the stellar atmosphere (where the models may be perfectly adequate) than the O I lines would then be incorrect. However, the spectroscopic temperatures determined here are derived from such weaker, low-excitation lines. Additionally, the evaluation of the CFP temperature scale as being correct was also based on weaker, low-excitation lines. These facts suggest that the O I 7774 Å discrepancy is truly caused by an inaccurate temperature scale and that altering that scale is more than an “effective” way of providing a solution to the discrepancy.

The resolution to the 7774 Å O I controversy proposed here is, no doubt, a controversial one. However, one should consider these three *independent* findings: a) The temperature scale of CFP appears to be correct based on the spectroscopic constraints noted by Sneden *et al.* (1992); this scale is ~ 160 K hotter than that of M87 and C83 b) The recent spectroscopic fine analysis of HD 103095 by Smith *et al.* (1992) indicates that this subdwarf’s temperature is ~ 170 hotter than the photometric value based on the relation of C83 c) the color- T_{eff} relations for metal-poor stars derived here yield temperatures ~ 170 K hotter than the relations of M87 and C83. These independent results are remarkably consistent and suggest that photometrically derived temperatures for metal-poor dwarfs are typically 150 – 200 K too low. Fine analyses (like those of Smith *et al.*, 1992 and Sneden *et al.*, 1992) using high resolution and high S/N spectroscopy of only moderately metal-poor ($[\text{Fe}/\text{H}] = -1.0$) dwarfs can test this hypothesis.

Table 1. Comparison of 7774 Å O I Triplet Equivalent Widths

Star	EW(AR) mÅ	EW(TLLS) mÅ	EW(FOS) mÅ	EW(PAL) mÅ	EW(SLW) mÅ
BD+72 0094	87	74.3	—	—	—
HD 19445	39	31.5	—	—	29
HD 64090	>26	>17.8	—	—	—
HD 84937	76	39.8	—	—	—
BD+34 2476	65	61.5	—	—	—
HD 134169	166	—	152.0	—	150
HD 140283	>11	>12.6	—	>14	>15
BD+18 3423	154	—	138.7	—	—
BD+02 3375	52	32.5	—	—	—
HD 160693	203	—	168.3	169.3	—
BD+20 3603	43	35.1	34	—	—
BD+26 3578	53	39.7	—	—	—
HD 184499	203	—	—	189.6	—
HD 193901	96	80.0	—	—	—
HD 195633	223	—	227.5	204.0	—
BD+04 4551	108	97.6	—	—	—
HD 201891	135	—	119.4	120.7	—
HD 208906	205	—	—	144.2	—
HD 219617	87	—	—	70.2	—

Table 2. Photometric Data for Metal-Poor Stars

Star	[Fe/H]	$(V - K)_o$	$(B - V)_o$	$(b - y)_o$
HD 7424	-0.91	1.64	—	0.423
HD 16031	-1.91	1.31	0.44	0.323
HD 19445	-2.15	1.39	0.46	0.349
HD 74000	-2.06	1.26	0.42	0.317
HD 76932	-0.98	1.45	0.51	0.364
HD 84937	-2.24	1.21	0.39	0.303
HD 94028	-1.66	1.39	0.48	0.343
HD 106516	-0.84	1.245	—	0.317
HD 108177	-1.87	1.32	0.43	0.323
HD 132475	-1.93	1.625	0.555	0.386
HD 134169	-0.96	1.46	—	0.376
HD 140283	-2.64	1.580	0.48	0.377
HD 148816	-0.74	1.42	—	0.362
HD 184499	-0.51	1.34	—	0.390
HD 188510	-1.60	1.69	0.59	0.408
HD 193901	-1.22	1.52	0.55	0.371
HD 194598	-1.32	1.380	0.485	0.351
HD 201891	-1.08	1.420	0.510	0.348
HD 219617	-1.63	1.39	0.48	0.342
BD -4 3208	-2.53	1.28	0.395	0.304
BD +2 3375	-2.40	1.404	0.43	0.352
BD +2 4651	-1.98	1.287	0.40	0.318
BD +17 4708	-1.98	1.395	0.44	0.325
BD +21 607	-1.89	1.325	0.45	0.321
BD +26 2606	-2.63	1.37	0.425	0.328
BD +29 2091	-2.20	1.50	0.50	0.377
BD +34 2476	-2.19	1.23	0.395	0.301
BD +36 2165	-1.74	1.27	0.43	0.314
BD +51 1696	-1.49	1.60	0.55	0.392
BD +54 1216	-1.75	1.335	0.48	0.336

Table 3. Spectroscopic T_{eff} Values from Line Ratios

Star	T_{eff} (6748/6750)	T_{eff} (6748/6752)	T_{eff} (6297/6290)	T_{M} ($B - V$)	T_{M} ($b - y$)	T_{M} ($V - K$)	T_{C} ($V - K$)
BD +18 3423	6164	—	—	5843	5902	—	—
	6166	6172	—	5843	5902	—	—
HD 184499	—	—	5926	—	5589	5983	5984
HD 189558	5919	6042	—	5687	5615	—	—
HD 201891	—	—	6141	5726	5869	5859	5848
HD 205650	5933	6125	—	5726	5702	—	—

Table 4. Comparison of Spectroscopic and Photometric Temperatures I.

Star	T(Spectra)	$T(B - V)$	$T(b - y)$	$T(V - K)$	$T(V - K)$
	Perrin <i>et al.</i>	Magain	Magain	Magain	Carney
HD 19445	5929	5921	5862	5906	5898
HD 25329	4898	—	—	—	4705
HD 64090	5358	5335	5322	—	5360
HD 94028	5998	5843	5902	5906	5898
HD 103095	5140	—	—	—	4996
HD 132475	5861	5550	5615	5541	5525
HD 140283	5929	5843	5675	5611	5594
HD 219617	5861	5843	5908	5906	5898

Table 5. Comparison of Spectroscopic and Photometric Temperatures II.

Star	T_{sp} K	$T_{\text{CFP}}(V - K)$ K	$T_{\text{C}}(V - K)$ K
HD 81192	4750	4760	4605
HD 88609	4500	4710	4535
HD 122563	4600	4650	4465
HD 184266	6000	—	5360
BD +58 1218	5000	5150	5020
BD -18 5550	4750	4750	4595

Table 6. Reanalysis of [O/Fe] Ratios in Metal-Poor Stars

Star		T _{eff} K	[Fe/H]	Refs [†]	[O/Fe] LTE	[O/Fe] NLTE
HD	7424	5639	-0.73	b,c	+0.60	+0.58
HD	16031	6180	-1.78	f,b,c,d	+0.48	+0.46
HD	19445	6007	-2.06	f,b,d	+0.64	+0.64
HD	64090	5523	-1.66	f,b,c	+0.56	+0.56
HD	74000	6220	-2.02	f,b,c,d	+0.38	+0.36
HD	84937	6314	-2.21	f,b,c	+0.44	+0.40
HD	87140	5133	-1.60	b,g,h,i	+0.51	+0.49
HD	103095	5170	-1.30	a	+0.43	+0.43
HD	188510	5613	-1.52	b,c	+0.71	+0.71
HD	201891	6013	-1.00	c,d,e	+0.42	—
BD	-4 3208	6307	-2.43	b,d	+0.56	+0.48
BD	+2 3375	5987	-2.35	f,b,c	+0.71	+0.69
BD	+3 740	6254	-2.76	f,b	+0.34	+0.26
BD	+37 1458	5425	-2.11	b,c,g	+0.65	+0.63
BD	+29 2091	5820	-1.91	b,d,g,j	+0.47	+0.46

[†] *References:* a) Smith *et al.* 1992 b) TLLS c) Laird *et al.* 1988
d) Ryan & Norris 1991b e) Lambert *et al.* 1991 f) Magain 1987b,
Magain 1989 g) Tomkin, Sneden & Lambert 1986 h) Gratton
& Sneden 1987 i) Schuster & Nissen 1989 j) Rebolo, Molaro &
Beckman 1988

Figure 1. $(b - y) - T_{\text{eff}}$ relations (for $\log g = 4.0$) from new Kurucz (1991) models. The different curves correspond to stellar models having differing metallicities. The curves, from top-right to bottom-left, represent models with $[\text{Fe}/\text{H}] = +0.2, +0.0, -0.3, -0.5$, *etc.* . At $(b - y) \leq 0.42$, the curves for $-2.5 \leq [\text{Fe}/\text{H}] \leq -0.5$ essentially run together and form a roughly singular relation.

Figure 2. The $(b-y) - T_{\text{eff}}$ relation for $[\text{Fe}/\text{H}] = -0.5$ of SH, which should be valid for all halo stars in the range of color shown, compared to the analogous metallicity-independent relation of M87 for halo stars. The M87 relation yield values of T_{eff} which are ~ 150 K less than the SH relation.

Figure 3. The $(B - V) - T_{\text{eff}}$ relations (for $\log g = 4.0$) from new Kurucz (1991) models. The different curves correspond to stellar models having differing metallicities. The curves, from top-right to bottom-left, represent models with $[\text{Fe}/\text{H}] = +0.2, +0.0, -0.3, -0.5, -1.0, -1.5, -2.0, -2.5$. Even at the lowest $[\text{Fe}/\text{H}]$ values, there is still significant sensitivity of the inferred temperature to metallicity.

Figure 4. The $(B - V) - T_{\text{eff}}$ relations from the Kurucz models modified to be on the SH scale as described in the text compared to the metallicity-independent relation of M87 (solid line). The short dashed curve is the modified Kurucz relation for $[\text{Fe}/\text{H}] = -1.0$ while the long dashed curve is that for $[\text{Fe}/\text{H}] = -2.5$. For the more metal-rich stars, the temperatures derived from the M87 relation are ≥ 200 K lower than those derived from the modified Kurucz relations.

Figure 5. The $(V - K) - T_{\text{eff}}$ relations (for $\log g = 4.0$) from new Kurucz (1991) models. The different curves correspond to stellar models having differing metallicities. The curves, from bottom-left to top-right, represent models with $[\text{Fe}/\text{H}] = +0.2, +0.0, -0.3, -0.5, -1.0, -1.5, -2.0$, and -2.5 . Except for the lowest metallicities the models imply there is some small sensitivity of the inferred temperature to metallicity.

Figure 6. The temperature difference $T(b - y) - T(B - V)$ calculated from the M87 calibrations *vs.* $[Fe/H]$. The trend seen is statistically significant and indicates that there appear to be true metallicity dependences in color- T_{eff} relations which are not taken into account by M87. The trend seen is in the same sense predicted by the Kurucz models.

Figure 7. a) The temperature difference $T(b-y) - T(V-K)$ calculated from the M87 calibrations *vs.* $[Fe/H]$. While there is perhaps a suggestion of a trend in the same sense as predicted by the Kurucz models, the significance level is marginal (slightly less than 2σ) and sensitive to the statistical test used. b) Same as Figure 7a but with additional points added to increase the metallicity baseline. The trend seen here is statistically significant (even if one excludes the very discrepant, most metal-rich point) and in the sense predicted by the Kurucz models.

Figure 8. The $(V - K) - T_{\text{eff}}$ relation for Pleiads in C83 which have colors in the valid range of the SH calibration.

Figure 9. $[O/Fe]$ *vs.* $[Fe/H]$ for the TLLS subsample reanalyzed here. The mean abundance ratio is $[O/Fe] = +0.53$ ($\sigma = 0.12$ dex). There is no statistically significant trend of $[O/Fe]$ with $[Fe/H]$ in this sample. The $[O/Fe]$ ratios are in agreement with those found in giants using the 6300 Å [O I] line.

Figure 10. a) $[O/Fe]$ ratios of TLLS versus their adopted T_{eff} values for those stars in the subsample reanalyzed here. The correlation between $[O/Fe]$ and T_{eff} is significant at a very high confidence level (as TLLS found for their complete sample). b) $[O/Fe]$ *vs.* T_{eff} for the TLLS subsample reanalyzed here. No statistically significant correlation is found in the reanalysis.

Chapter 3. Possible Cosmological Implications of a Hotter Temperature Scale for Halo Stars

Abstract

Possible cosmological implications of a hotter T_{eff} scale for field halo stars are addressed. Additional evidence, including a recent spectroscopic T_{eff} determination for the halo star HD 89499; T_{eff} values derived from Balmer line modeling for two halo stars; and temperatures derived using transformations from two independent sets of widely-used isochrones, is presented which supports the need for such an upward revision in the T_{eff} scale. Additionally, we note that the hotter temperature scale is consistent with the Bell & Gustafsson (1989) relations. Temperatures inferred from both the Victoria and Yale isochrone color- T_{eff} conversions are in good agreement with the revised T_{eff} scale for field stars on the upper main sequence and subgiant branch. We conclude that no downward revision of globular cluster turnoff ages would result for those clusters having $[\text{Fe}/\text{H}]$ near -1.3 . A similarly strong conclusion can not be reached for clusters of other metallicities, but it seems likely that no substantial age reductions ($\gtrsim 1$ Gyr) are necessary due to uncertainties in color- T_{eff} relations. Although large uncertainties in low mass stellar models are present, the good agreement between the new Yale isochrones, which are redder at cooler temperatures than the Victoria isochrones, and the observed lower main sequence of M92 seems to be consistent with the somewhat larger upward revisions for the temperatures of cooler field stars that has been advocated. The high T_{eff} scale leads to a mean ${}^7\text{Li}$ abundance of 2.28 ± 0.09 (1 standard deviation) in hotter halo stars. Then, a conservative minimal primordial value of $N(\text{Li}) = 2.35$ is inferred. When this value is combined with recent estimates of the primordial ${}^4\text{He}$ abundance, not taking into account the observational and model errors, and the limit set by the primordial $\text{D}+{}^3\text{He}$ abundance, inconsistency with standard Big Bang nucleosynthesis models occurs. However, if the observational *and* theoretical uncertainties are included, consistency with the standard model is still possible in the range $0.012 \leq \Omega_b \leq 0.017$ (for $H_0 = 100$ km/s/Mpc). The higher T_{eff} scale raises the observed B/Be ratio in HD 140283 from 10 to 12. This new value is more comfortably consistent with production of the observed B and Be by ordinary spallation processes, but does not preclude a significant primordial component in more metal-poor stars.

1. Introduction

By remarking that "... as long as galaxies are made of stars, our understanding, our system

of datation, our survey of the depths of space depend on our knowledge of stars”, E. Schatzman (Brown *et al.* 1993) reminds us that much of our current understanding of cosmology hinges on our knowledge of fundamental *stellar* parameters. Depending on one’s point of view, this may or may not be a fortunate of circumstance. On one hand, this means that further and better understanding of cosmological issues “only” requires associated knowledge provided by the study of relatively bright, nearby stars. On the other hand, it is perhaps frustrating that conclusions regarding big bang nucleosynthesis or the age of the universe are held hostage by the uncertainty in stellar temperatures, gravities, and composition or by the lack of detailed knowledge of stellar physics (*e.g.*, elemental diffusion, rotational mixing, color- T_{eff} relations).

Chapter 2 suggested that the current halo star temperature scale (as defined by color- T_{eff} relations of, *e.g.*, Magain 1987a or Carney 1983) may be ~ 150 K too low. This conclusion was based on investigations of three sets of stellar temperatures: a) the T_{eff} for Gmb 1830 derived by a recent spectroscopic fine analysis b) the T_{eff} scale of Cohen, Frogel & Persson (1978), whose accuracy was demonstrated by the lack of correlation between derived metal abundances and excitation potential when their T_{eff} values are assumed and c) the $(b - y)$ - and $(V - K) - T_{\text{eff}}$ relations for halo stars derived in K93 by tying the metal-poor T_{eff} scale to the Pop I T_{eff} scale of Saxner & Hammarback (1985). These three sets of temperatures are consistently $150 - 200$ K higher than the T_{eff} scale defined by the Magain (1987a) and Carney (1983) relations. Additionally, the three sets of temperatures can be shown to be mutually consistent, which argues for their accuracy. We also showed that other, probably less precise, spectroscopic T_{eff} estimates for halo stars tend to be larger than current photometric estimates. If one adopts the temperatures from the revised $(b - y)$ and $(V - K)$ relations, the [O/Fe] ratio (as derived from the O I 7774 Å triplet) of halo dwarfs is in accord with [O/Fe] ratio observed in halo giants from the 6300 Å [O I] line. Also, the trend between [O/Fe] and T_{eff} seen by Tomkin *et al.* (1992) is removed.

Such a revision in the halo star T_{eff} scale, if truly warranted, could have serious implications for cosmology. Since these were not discussed in Chapter 2, we take the opportunity to do so now. The primordial abundance of the light elements, especially ${}^7\text{Li}$, can provide constraints on cosmology via comparison with models of Big Bang nucleosynthesis (*e.g.*, Boesgaard & Steigman 1985, Deliyannis *et al.* 1989). The primordial ${}^7\text{Li}$ abundance is inferred from its abundance in halo stars. As is well-known, this abundance is sensitive to the adopted T_{eff} of the star. The B/Be ratio in halo stars (Duncan, Lambert & Lemke 1992) can potentially serve to discriminate between homogeneous and inhomogeneous models of Big Bang nucleosynthesis. The B abundance in halo stars (as derived from the neutral resonance lines at 2497 Å) is also sensitive to the adopted temperature. Globular

cluster isochrones (*e.g.*, Bergbusch & Vandenberg 1992), used to provide absolute ages for old stellar systems, could also be affected by changes in the halo star T_{eff} scale— both in the fixing of the parameter α and in the color— T_{eff} relations used to transform the theoretical calculations to the observational plane (Vandenberg 1992). Because globular clusters provide an important lower limit to the cosmic age, any changes in their ages resulting from revisions in halo star T_{eff} values are of interest. Specifically, raising the halo star T_{eff} scale by ~ 150 K could lower globular cluster ages by a couple of Gyr *if* this raising also implies an associated revision in the cluster distance modulus (*i.e.*, alters the turnoff luminosity). We stress that isochrones’ predicted T_{eff} values are sensitive to various physical assumptions included (or excluded) in the models; therefore, it is to be emphasized that relying on T_{eff} *per se* to determine cluster ages is something we do not advocate here. In any case, *if* such adjustments to the T_{eff} scale do also imply revised distances, then such an effect might place the age of M92 at ~ 12 Gyr (based on the 14 Gyr assignment of Bergbusch & Vandenberg 1992). Note that a 12 Gyr age is just compatible with recently measured large values (~ 80 km/s/Mpc) of H_0 in a flat Universe.

2. Further Evidence for a Revised Halo Star T_{eff} Scale

2.1 Comparison with T_{eff} scale of Bell & Gustafsson

Comparison of the new temperatures in chapter 2 with those of Bell & Gustafsson (1989) is of interest. Bell & Gustafsson (1989) used the results of infrared surface flux analysis and spectral synthesis calculations to determine color— T_{eff} relations for G and K stars. We compared a number of the T_{eff} values of stars from Table 6 of chapter 2 and Table 7 of this chapter with the T_{eff} values derived from the Bell & Gustafsson (1989) relations. For the warmer stars ($T_{\text{eff}} \gtrsim 5400$ K) we checked, the agreement was excellent. In fact, the Bell & Gustafsson relations tended to give slightly higher temperatures (by ~ 15 K) and the dispersion was small (± 35 K). For cooler stars, their temperatures seem significantly higher. For HD 103095 a value of 5190 K was found compared to 5115 K on the chapter 2 scale; and for HD 87140, a value of 6000 K was determined compared to 5133 K on the chapter scale. The Bell & Gustafsson results seem to reinforce the conclusion that larger upward revisions of the effective temperatures are needed for cooler stars. No comparison has been made with the temperatures from the relation of Ridgway *et al.* (1980) since this is for stars much cooler than those we are interested in here. However, Bell & Gustafsson (1989) do note that their T_{eff} values are typically 100 K hotter than those of Ridgway *et al.* (1984).

2.2 Halo Star Temperatures from Balmer Line Profiles

As is well known, the Balmer line profiles in F and G stars are very temperature sensitive. Thus, the profiles serve as good temperature indicators. Fuhrmann, Axer & Gehren (1993) have discussed the modeling of the Balmer line profiles as a means to determine T_{eff} and $\log g$ for F and G stars of all metallicities. They determine temperatures for two halo stars, HD 140283 and G41-41. The result is $T_{\text{eff}} = 5810$ K for HD 140283. Also, consistency between the $H\alpha$ and $H\beta$ results requires that this star be a subgiant ($\log g \sim 3.3$), an important finding as the gravity of this critical subdwarf has been in question. Their T_{eff} estimate is in close agreement with the value derived on the scale of chapter 2, 5813 K (see Table 7). However, this value is ~ 200 K hotter than the $(V - K)$ photometric estimates from the relations of Carney (1983) and Magain (1987a) and is ~ 135 K hotter than the $(b - y)$ photometric value from the relation of Magain (1987a). As a point of interest, we note that the Bell & Gustafsson (1989) relations give $T_{\text{eff}} = 5750$ K.

Fuhrmann *et al.* (1993) find $T_{\text{eff}} = 6450$ for G41-41. The $(V - K)$ and $(b - y)$ indices of this star are available in Carney (1983) and Hauck & Mermilliod (1990). A reddening of $E(B - V) = 0.01$ is given by Laird, Carney & Latham (1988). Using the color- T_{eff} and adopted reddening relations in chapter 2, temperatures of 6340 and 6361 K are found from $(b - y)$ and $(V - K)$. These values may be compared with 6160 and 6134 K derived from the Carney (1983) and Magain (1987a) $(V - K)$ relations and 6196 K from the Magain (1987a) $(b - y)$ relation. For those interested, the T_{eff} from the Bell & Gustafsson (1989) relation is ~ 6500 K, but this value is uncertain as colors for 6500 K are given for only one value of $\log g$. Again, we see that the chapter 2 values are hotter than those on the Magain (1987a) and Carney (1983) scale and are in closer agreement with the spectroscopic estimate from the Balmer line profiles. The offset between the Fuhrmann *et al.* (1993) value and the chapter 2 values could be due to an underestimate in the reddening of G41-41. If $E(B - V) = 0.03$, then excellent agreement exists between the K93 values and the Balmer line value. Indeed, the reddening derived from Stromgren photometry (Schuster & Nissen 1988) is completely consistent with $E(B - V) = 0.03$. In any case, we believe these comparisons provide further evidence that the temperatures derived from the Carney (1983) and Magain (1987) relations, commonly used in stellar abundance studies, are too low for halo stars.

2.3 Spectroscopic T_{eff} of HD 89499

Balachandran *et al.* (1993) have recently reported the detection of ^7Li in the cool halo binary HD 89499. Their analysis provides one of the few modern spectroscopic T_{eff} determinations for a

halo star. Using the $(V - K)$ index and the $(V - K) - T_{\text{eff}}$ relation of Carney (1983), they find $T_{\text{eff}} = 4814$ K. They then compare derived abundances from 26 Fe I lines against the lines' excitation potential. The oscillator strengths used here are “astrophysical” ones, found by determining the $\log gf$ values which reproduce an assumed Fe abundance in the Sun using the solar line strengths.

Figure 11a shows Balachandran *et al.*'s abundances derived in this way *vs.* excitation potential (where the data has been taken directly from their Table 1). The dashed line shows the mean $[\text{Fe}/\text{H}]$. As Balachandran *et al.* state, there is no apparent correlation. Indeed, the ordinary correlation coefficient in this case is $r = 0.022$, indicating any slight slope is significant at the $< 50\%$ confidence level. The essentially zero slope indicates that the temperature assumed, $T = 4850$ K, is reasonably correct *given the choice in oscillator strengths*. However, as Balachandran *et al.* (1993) note, the Fe I lines used in the study of HD 89499 ($[\text{Fe}/\text{H}] \sim -2$) are strong in the Sun. Hence, the deduced oscillator strengths will be sensitive to the adopted solar microturbulent velocity and the treatment of damping. Additionally, because many of the lines will lie on the flat portion of the curve of growth, any errors in equivalent width will be magnified in the $\log gf$ determinations.

Balachandran *et al.* (1993) found that use of experimental oscillator strengths resulted in a T_{eff} value of 4950 K. An intermediate value of 4900 K was adopted as a final compromise. It is interesting to see what the implications of theoretical oscillator strength calculations are for the T_{eff} of this star. Dr. R. Kurucz kindly provided his theoretical $\log gf$ values for the 24 Fe I lines with $\lambda > 5000$ Å. The value for the 7780.57 Å line is very discrepant with the solar gf value given by Balachandran *et al.* (1993) and the value given by Gurtovenko & Kostik (1981) and we excluded this line from further consideration. Abundances were calculated for each line by changing the abundance in Table 1 of Balachandran *et al.* (1993) by the amount corresponding to the difference in their solar gf values and the Kurucz gf values (since, in LTE analyses, the quantity $\log gfN$ is, all other things being equal, a constant). A plot of the derived abundances versus the excitation potential is shown in Figure 11b. The trend in the data is marked. The correlation coefficient is 0.740, indicating a statistically significant correlation at the $\geq 99.9\%$ confidence level.

The direction of the trend is such that it requires a *higher* T_{eff} value than assumed (4850 K). We suggest that the results of the experimental and theoretical $\log gf$ values argue for the hotter T_{eff} (4950 K) found by Balachandran *et al.* (1993). The temperature derived from the $(V - K)$ index using the T_{eff} scale of K93 is 4935 K, in excellent agreement with the higher, spectroscopic value and ~ 120 K larger than the value derived from the $(V - K) - T_{\text{eff}}$ relation of Carney (1983).

3. The Primordial ${}^7\text{Li}$ Abundance

The primordial ${}^7\text{Li}$ abundance provides strong constraints on models of Big Bang nucleosynthesis (*e.g.*, Boesgaard & Steigman 1985 and Deliyannis *et al.* 1989). The pioneering work of Spite & Spite (1982), seemingly confirmed by subsequent observations, has placed the primordial Li abundance at $N(\text{Li}) \sim 2.1$ as determined from abundances in the atmospheres of metal-poor halo stars with the assumption that no depletion has occurred. Because the ${}^7\text{Li}$ I line at 6707 Å is fairly temperature sensitive, any change in the T_{eff} scale of halo stars requires a concomitant change in the derived ${}^7\text{Li}$ abundances.

This change has been investigated using a selection of halo stars with ${}^7\text{Li}$ abundance measurements in the literature. Particular attention was paid to the selection of hotter, more metal-poor stars to avoid those stars which have possibly suffered from substantial depletion. Stars were selected mainly from the work of Hobbs & Thorburn (1991); Rebolo, Molaro & Beckman (1988); and Spite & Spite (1982). Temperatures were derived using the $(b - y) - T_{\text{eff}}$ relation given in K93. The values for $(b - y)$ were taken from the catalogue of Hauck & Mermilliod (1990). Dereddened indices were constructed using the $E(B - V)$ values mainly from Laird, Carney & Latham (1988) via the relation $E(b - y) = 0.72E(B - V)$. Abundances in the literature were then adjusted for the new value of T_{eff} given the sensitivity of abundance to T_{eff} listed in the reference source. The first column of Table 7 lists the stars reanalyzed in this fashion. Column 2 lists the $[\text{Fe}/\text{H}]$ value calculated from various sources in the literature after altering the abundance for the specific T_{eff} adopted here (again, according to the T_{eff} sensitivity given in the source) and adjusting for a homogeneous solar iron abundance $\log N(\text{Fe}) = 7.51$ (Biemont *et al.* 1991; Holweger *et al.* 1991; Holweger, Heise & Kock 1990). The main sources used for the $[\text{Fe}/\text{H}]$ data were Magain (1987b, 1989); Tomkin *et al.* (1992); Gratton & Sneden (1987); Tomkin, Sneden & Lambert (1986); Laird *et al.* (1988); Ryan & Norris (1991); and Beers *et al.* (1990). Columns 3 and 4 of Table 7 give our adopted T_{eff} values and new values for $N(\text{Li})$.

The mean abundance³ is $N(\text{Li}) = 2.28 \pm 0.09$ dex. This uncertainty is a standard deviation (not an error in the mean) and seems consistent with the uncertainty in the equivalent widths, adopted T_{eff} , and other parameters (however, for a different point of view see Deliyannis, Pinsonneault & Duncan 1993). This reanalysis was performed before the recent results of Thorburn & Beers (1993) were published. Their new equivalent width measurements for G 64-12 and BD+3 740 are $\sim 16\%$

³ Calculated excluding HD 134169, which has a rather uncertain metallicity and so may not be a genuine member of the halo but rather a disk (or thick disk ?) star with depleted Pop I ${}^7\text{Li}$.

and $\sim 10\%$ higher than the values given in Hobbs & Thorburn (1991). In the interest of being conservative, we adopt the lower values. Although the two new measurements alone hardly provide any evidence of systematic errors in previous equivalent width estimates, higher resolution and/or S/N observations may be desirable to exclude the existence of such an effect.

4. The Primordial B/Be Ratio

4.1 The Observed B/Be Ratio in HD 140283

The primordial B/Be ratio also can potentially provide constraints on models of Big Bang nucleosynthesis and Galactic chemical evolution. Duncan, Lambert & Lemke (1992) have utilized HST observations of the B I resonance lines at $2497 \text{ \AA}$ to estimate the B/Be ratio of the very metal-poor halo star HD 140283. Their estimated ratio for this important object is $\text{B/Be} = 10^{+5}_{-4}$, which is also the lower limit predicted by models of synthesis of Be and B via cosmic ray spallation reactions in the ISM (see below).

Like the $6707 \text{ \AA}$ Li line, the B I resonance line is sensitive to temperature. Any change in the adopted T_{eff} results in a change in the derived B abundance and the B/Be ratio. Using the revised T_{eff} given in Table 7 (5813 K), the adopted T_{eff} (5640 K) of Duncan *et al.*, and the sensitivity they give ($\Delta \log \epsilon(\text{B})/\epsilon(\text{Be}) = +0.05 \text{ dex}$ for $\Delta T_{\text{eff}} = +100 \text{ K}$), the B/Be ratio would *rise* from 10 to 12. This value seems to be even more consistent with production of the observed B and Be by spallation reactions rather than indicating significant primordial contributions (see the predictions of, *e.g.*, Kawano *et al.* 1991).

4.2 Predicted B/Be Ratios from Spallation Models

As noted above, Duncan *et al.* (1992) find that the minimum B/Be ratio produced by cosmic ray spallation reactions is 10. Their spallation model assumes $[\text{O/Fe}] = +0.5$, $[\text{C/Fe}] = 0$, and $[\text{N/Fe}] = 0$ for $[\text{Fe/H}] < -1.0$ halo dwarfs based on the discussion in Lambert (1989). This argument is not entirely consistent however. In fact, without a revision of the halo star T_{eff} scale, no study to date has indicated that $[\text{O/Fe}]$ is as low as $+0.5$ in halo dwarfs (or mild subgiants) having the metallicity of HD 140283. The results of Tomkin *et al.* (1992) yield $[\text{O/Fe}] \sim +0.8$ at the metallicity of HD 140283. The surprising results of Abia & Rebolo (1988) indicate $[\text{O/Fe}] = 1.0 - 1.1$ at the metallicity of HD 140283. Bessell, Sutherland & Ruan (1991) find $[\text{O/Fe}] = +0.65$ and $+0.9$ for HD 140283 (assuming the same $[\text{Fe/H}]$ as Duncan *et al.* do). Sneden, Lambert & Whitaker (1979) find $[\text{O/Fe}] = +0.95$ (again, assuming the same $[\text{Fe/H}]$ as Duncan *et al.* do) for HD 140283. Because the adopted $[\text{O/Fe}]$

ratio determines the C/O and N/O ratios (given assumptions about [C/Fe] and [N/Fe]) to be used in the B/Be spallation models, we emphasize that $[\text{O}/\text{Fe}] = +0.5$ can only be obtained for HD 140283 (and dwarfs of similar metallicity) by adopting a T_{eff} scale similar to that proposed by K93. Indeed, the T_{eff} and $[\text{Fe}/\text{H}]$ values for HD 140283 given in Table 7 and the data of Tomkin *et al.* (1992) fortuitously result in an LTE abundance ratio of exactly $[\text{O}/\text{Fe}] = +0.50$.

Fortunately, as Duncan *et al.* (1992) point out, the spallation model B/Be ratio is fairly insensitive to the $[\text{O}/\text{Fe}]$ ratio. In order to confirm this, we calculated (see Duncan *et al.* for computational details) spallation production ratios using the Read & Viola (1984) cross sections, $[\text{O}/\text{Fe}]$ ratios of both +0.5 and +1.0, abundance ratios $[\text{C}/\text{Fe}] = [\text{N}/\text{Fe}] = 0$, and the solar CNO abundances as given by Grevesse & Anders (1989). The predicted B/Be ratio for $[\text{O}/\text{Fe}] = +0.5$ is 10.4. The predicted B/Be ratio for $[\text{O}/\text{Fe}] = +1.0$ is 9.9. Hence, changing the O/Fe ratio by a factor of ~ 3 only results in a $\sim 5\%$ change in the predicted B/Be ratio. This result will be of particular importance when B/Be ratios are obtained for stars even more metal-poor than HD 140283. The recent models of both Prantzos *et al.* (1993) and Matteucci & Francois (1992) indicate that $[\text{O}/\text{Fe}]$ may rapidly increase for $[\text{Fe}/\text{H}] \leq -3$. Thus, even without knowing the $[\text{O}/\text{Fe}]$ ratio in stars having $[\text{Fe}/\text{H}] \leq -3$, one may develop relatively secure models for comparison with the observed B/Be ratios.

5. Implications for Big Bang Nucleosynthesis

5.1 The Primordial ${}^7\text{Li}$ Abundance

Before comparison can be made with models of Big Bang nucleosynthesis, the revised Li plateau abundance ($N(\text{Li}) = 2.28$) found here must be related to the true primordial ${}^7\text{Li}$ abundance, which we take here as the *initial* halo star abundance. Computations which yield such initial abundances after following the evolution of surface ${}^7\text{Li}$ abundance in halo stars from standard stellar evolution models have been discussed by Deliyannis, Demarque & Kawaler (1990). We focus on the results from standard stellar evolution models for two reasons: 1) They yield more conservative (*i.e.*, lower) estimates for the primordial Li abundance than models that include, *e.g.*, diffusion or rotation and 2) An upward revision of the halo star T_{eff} scale of the magnitude assumed here would seemingly eliminate the efficacy of (non-standard) diffusion mechanisms in depleting the initial Li abundance of halo dwarfs (C. Deliyannis, private communication 1993).

Deliyannis *et al.* (1990) find, by fitting their extensive Li isochrones to the available data, that the observed ${}^7\text{Li}$ plateau (on the “old”, cooler T_{eff} scale) at $N(\text{Li}) = 2.08$ corresponds to a primordial Li abundance of 2.17. From a) the shift of the data points in figure 27 of Deliyannis *et al.* (1990)

required by the higher T_{eff} scale used here and b) the fits listed in their Table 13, one may infer that the adjustment which yields the primordial abundance is *larger* than 0.09 dex as determined by Deliyannis *et al.* for the T_{eff} scale they use. In the interest of being conservative, however, we will adopt a slightly *smaller* value. Based on our observations and a conservative estimate of stellar Li depletion from standard stellar evolution models, we adopt a primordial Li abundance of $N(\text{Li}) = 2.35 \pm 0.09$ (where, for the sake of being conservative, the uncertainty is 1 standard deviation rather than our error in the mean).

5.2 Comparison with Big Bang Nucleosynthesis Models

We now compare light element abundances with the detailed model results of Krauss & Romanelli (1990). Using a Monte Carlo analysis, they provide confidence levels for standard Big Bang model predictions by careful consideration of the uncertainties in the various reaction rates and the neutron half-life. We use Krauss & Romanelli's (1990) figures 17 and 18 to compare with our primordial ${}^7\text{Li}$ abundance estimate of 2.35. The situation is also shown in Figures 12a and 12b, which graph Krauss & Romanelli's (1990) predicted primordial ${}^4\text{He}$ and ${}^7\text{Li}$ abundances versus the logarithm of η (multiplied by 10^9), the usual nucleon-to-photon ratio (see *e.g.*, Boesgaard & Steigman 1985). Also shown in the figure are the primordial ${}^7\text{Li}$ abundance found here (the horizontal dashed line in Figure 12b), the minimum value of η implied by the maximum $\text{D}+{}^3\text{He}$ abundance of Boesgaard & Steigman (1985), and the value of η implied by a primordial ${}^4\text{He}$ value of $Y = 0.22$ (Fuller, Boyd & Kalen 1991). These values of η are shown by the solid vertical lines in Figures 12a and 12b.

If one adopts the recent value (*i.e.*, the value itself, making no consideration of the related uncertainties) of Fuller *et al.* (1991) for the primordial ${}^4\text{He}$ abundance, $Y_p = 0.220 \pm 0.010$, and the range of the *upper* limit to the $\text{D}+{}^3\text{He}$ abundance given by Boesgaard & Steigman (1985)– $(\text{D}+{}^3\text{He})_p \leq 6 - 10 \times 10^{-5}$ – an inconsistency with the standard model becomes apparent. If the ${}^7\text{Li}$ abundance is placed on the predicted descending branch (its abundance behavior is *not* a monotonic function of Ω_b , see Figure 12b), it is within 1σ of the Fuller *et al.* (1991) ${}^4\text{He}$ abundance. However, in this case, both values (again, making no consideration of the errors in these quantities for the moment) are inconsistent with the limit set by the primordial $\text{D}+{}^3\text{He}$ abundance range.

This inconsistency is essentially that noted by Fuller *et al.* (1991). However, now taking into account the errors in Y_p , we differ with them in that the 2σ range of their Y_p estimate is consistent with the upper limit to the primordial $\text{D}+{}^3\text{He}$ abundance. Our new primordial Li abundance estimate is very important in this regard as consistency requires that the primordial Li abundance

be on the ascending branch of the abundance *vs.* Ω_b fiducial. Even if this is the case, the 2σ (including *both* the observed uncertainties and the model uncertainties from Krauss & Romanelli 1990) ${}^7\text{Li}$ abundance is too large to overlap with the 2σ value of Fuller *et al.*'s (1991) Y_p estimate (although it is consistent with the $\text{D}+{}^3\text{He}$ limit).

Can consistency with the standard model of Big Bang nucleosynthesis be achieved? Although the consistency of the standard model is in a somewhat precarious position given the ${}^7\text{Li}$ results presented here, we argue that it still remains tenable if all the uncertainties are fully stretched. As we noted above, consistency is not achieved between the Fuller *et al.*'s $Y_p + 2\sigma(Y_p)$ value and our ascending branch $N(\text{Li}) - 2\sigma(\text{Li})$ value. However, the 2σ value for Y_p only included the observational uncertainty and not the uncertainty in the Big Bang predictions as computed by Krauss & Romanelli (1990). If one takes the 2σ value of Fuller *et al.* (1991), $Y_p = 0.240$, and places it on the 1σ fiducial of the Krauss & Romanelli (1990) calculations, consistency can be achieved with the 2σ ${}^7\text{Li}$ observations on the ascending portion of the abundance *vs.* Ω_b relation. The small range of consistency where the 2σ (now σ includes contributions from both observation and the Krauss & Romanelli model) Y_p and $N(\text{Li})_p$ values overlap is given by $0.012 \leq \Omega_b \leq 0.017$ for $H_0 = 100$ km/s/Mpc. In any case, the upper 2σ value for ${}^7\text{Li}$ itself yields a limit $\Omega_b \leq 0.019$ (*i.e.*, the new Li abundances found here do not change the usual conclusion that if one desires $\Omega = 1$, then non-baryonic matter seems necessary).

We note two points concerning these conclusions. First, the above comparisons assumed 3 neutrino species. If $N_\nu \geq 3.5$ then the standard model would be ruled out by the primordial ${}^7\text{Li}$ abundance suggested here. We find that for $N_\nu \leq 3.5$ and the 2σ uncertainties of the Krauss & Romanelli (1990) models, one can not rule out the standard model on the basis of the Y_p of Fuller *et al.* (1991) and the $\text{D}+{}^3\text{He}$ limit when all uncertainties are considered. This illustrates the great utility of ${}^7\text{Li}$ in potentially constraining Big Bang nucleosynthesis. Secondly, the B/Be ratio of HD 140283 (either that of Duncan *et al.* 1992 or the slightly higher one found here) is completely consistent with production of these elements by ordinary cosmic ray spallation and does not indicate a significant primordial component. However, the result certainly does not preclude the possibility of primordial production in a non-standard Big Bang nucleosynthesis scenario. Such a conclusion will have to await further measurements of B and Be abundances in even more metal-poor halo stars.

6. Globular Cluster Ages

6.1 The Victoria Group's Isochrones

It has been repeatedly noted (*e.g.*, Vandenberg & Poll 1989) that the original isochrones of Vandenberg & Bell (1985) seem to require, on average, a color shift of 0.02-0.03 mag to attain the best agreement with the observed morphology of globular cluster HR diagrams (once best estimates of distance modulus and reddening have been selected)⁴. Indeed, the new O-enhanced isochrones of Bergbusch & Vandenberg (1992) have adopted this shift (see Vandenberg 1992) as it also results in better agreement with a reasonable solar color and the color of the calibrating subdwarf Gmb 1830. This redward color shift (analogous to a raising of the T_{eff} scale) results in better agreement with observations on three independent fronts and the revised T_{eff} scale that results is in close agreement with that suggested in chapter 2.

The Bergbusch & Vandenberg (1992, hereafter BV92) isochrones are normalized, in part, by matching observational parameters of the subdwarf Gmb 1830 (=HD 103095). By establishing the value of α (the usual mixing length to pressure scale height ratio, which is *assumed* to be independent of $[\text{Fe}/\text{H}]$, age, and evolutionary status), the zero point of the T_{eff} scale is fixed (Vandenberg 1992, hereafter V92). In this fixing of the temperature scale, V92 adopts $T_{\text{eff}} = 5105$ and $[\text{Fe}/\text{H}] = -1.36$. The metallicity is in good accord with the results of the fine analysis by Smith, Lambert & Ruck (1992) who find $T_{\text{eff}} = 5170$ for this star. The revised ($V - K$) calibration of chapter 2 yields 5115 K. Interestingly, the BV92 and revised Vandenberg & Bell (1985) T_{eff} scale is closer to the scale of chapter 2 than the scale established by the ($V - K$) relations of Carney (1983) or Magain (1987a).

Table 8 lists the stars in Table 6 of chapter 2 and Table 7 of this chapter which have $\log g$ values determined by Tomkin *et al.* (1992) and have T_{eff} values within a few hundred degrees of typical globular cluster turnoff stars. The columns give: (1) the star's identification (2) $[\text{Fe}/\text{H}]$ from either Table 6 Table 7 here (3) The $\log g$ value from Tomkin *et al.* (1992) (4) The T_{eff} from one of the two aforementioned tables (5) the T_{eff} that would be derived by interpolating within the V92 color- T_{eff} relations and (6) the difference between the two T_{eff} values- in the sense $T(\text{V92}) - T(\text{K93})$. The mean difference of the seven stars is 1.5 ± 13 K. We conclude that the Victoria isochrone T_{eff} scale is essentially the same as that suggested by chapter 2 for the point on the H-R diagram represented by HD 103095. The good agreement, noted above, of T_{eff} values for stars having different temperatures and metallicities than HD 103095 suggests that agreement may exist between the scale of chapter 2 and the Victoria scale in other locations. However, it must certainly be said that agreement is not guaranteed everywhere on all Victoria isochrones.

We conclude that halo star T_{eff} scale in chapter 2 would *not* lead to any revision in the absolute

⁴ Note that Demarque, Green & Guenther (1992) have suggested that such wholesale color shifts may not be a technically correct way to achieve such agreement.

turnoff ages of globular clusters having metallicities similar to HD 103095 as deduced from comparison with theoretical isochrones. Strictly speaking, we cannot draw such a conclusion for clusters of other metallicities as no other subdwarfs have been compared to the isochrone predictions. We can state that the color- T_{eff} relations used in the Victoria isochrones are in good agreement with those of chapter 2. However, since the Victoria color- T_{eff} relations and isochrones are not totally independent, the agreement between the Victoria T_{eff} values and those of chapter 2 for the stars in Table 8, which have temperatures and metallicities typical of globular cluster turnoff stars, would seem to argue against any significant shift in age. We envisage that in the most extreme case, the difference between V92's adopted temperature, 5105 K, for Gmb 1830 (or 5115 K derived from the $(V - K)$ calibration in chapter 2) and higher spectroscopic estimates (*e.g.*, 5140 K from Perrin *et al.* 1977; 5170 K from Smith, Lambert & Ruck 1992) might suggest a ~ 50 K shift of the Victoria isochrones to higher temperatures. This could produce an ~ 1 Gyr reduction at most in the age of a 15 Gyr cluster having $[\text{Fe}/\text{H}] \sim -1.3$.

6.2 The Revised Yale Isochrones

The other major set of isochrones often used in comparison with globular cluster observations is the *Revised Yale Isochrones* (Green, Demarque & King 1987). For the main sequence and subgiant stars, the new Yale transformations are derived in a way completely independent of how Vandenberg (1992) derives the new Victoria transformations. The only published information concerning the T_{eff} scale of these isochrones that we are aware of is that given by Green (1988, hereafter G88). From her Table 1, it is clear that, at metallicities and temperatures characteristic of globular cluster turnoff stars, the Yale isochrone $(B - V)$ colors are $0.02 - 0.03$ mag redder (*i.e.*, the T_{eff} values at a given $(B - V)$ color are hotter) than those of Vandenberg & Bell (1985). This means that the Revised Yale Isochrones T_{eff} values for globular cluster turnoff stars should be similar to the new O-enhanced Victoria isochrones and the isochrones of Vandenberg & Bell which have been shifted $0.02 - 0.03$ mag. This conclusion is confirmed in Table 8, which also gives the T_{eff} values and the differences from the calibrations of chapter 2 as derived from the Yale transformations included with the electronic version of the isochrones. The mean difference is 2.4 ± 51 K. Since the Yale isochrones and color- T_{eff} relations are independent, it cannot be concluded from the agreement between the temperatures advocated here and the RYI color conversion relation temperatures that the entire set of Yale Isochrones have a T_{eff} scale essentially identical to that of chapter 2. However, given the gross similarity between the Yale and Victoria isochrones, in that quite similar ages are derived for

globular clusters, it seems unlikely that any drastic reduction in globular cluster ages would result from the use of Yale isochrones either. Additionally, the fact that two different globular cluster isochrone color- T_{eff} transformation relations, derived completely independently from one another, agree with the T_{eff} scale suggested here seems to indicate the need for a higher T_{eff} scale for halo field stars.

6.3 Stars Cooler than the Turnoff

Table 1 of G88 indicates that color differences between the Yale Revised Isochrones and the Victoria isochrones are larger than 0.02 – 0.03 mag for the cool stars and, in fact, appear to be an increasing function of decreasing temperature⁵. Indeed, it was noted in chapter 2 that our revised T_{eff} scale for field halo stars seemed to indicate largest upward revisions were required for cooler stars. The Yale Revised Isochrone color conversion relation temperatures are expected *a priori* to be consistent with those of chapter 2 since the former are based (see description in §2.3 of G88) on the T_{eff} scale of Frogel, Persson & Cohen (1983). These latter temperatures are essentially those of Cohen, Frogel & Persson (1978), which were found to be consistent with ours as discussed in the previous chapter.

Do the larger color shifts at lower temperatures found by G88 mean that the T_{eff} values deduced from the new Victoria isochrones may be inconsistent with ours or G88's T_{eff} values at low temperatures ($T_{\text{eff}} \leq 5500$ K, say)? The answer would seem to be “no” for the red giant sequences. Just as the Revised Yale Isochrones force agreement with observed cluster RGB fiducials (see the oral discussion at the end of G88), the new O-enhanced isochrones of BV92 are forced (by arbitrarily altering surface pressure boundary conditions) to match the observed slopes of cluster RGB fiducials given by Frogel, Persson, & Cohen (1981). Again, these temperatures are on the scale of Cohen *et al.* (1978) which is consistent with that of chapter 2.

For the cool stars on the lower main sequence, large uncertainties exist in the isochrones. These are due to both uncertainties in the physics of low mass stars and possible deficiencies in model atmospheres. Additionally, photometry of globular cluster lower main-sequence stars is relatively uncertain due to both their faintness and the fact that various stellar photometry programs have differing treatments of sky subtraction which can yield different results (even for the same data). Unfortunately, these uncertainties are great enough so as to preclude any definite conclusions con-

⁵ Note that these increased offsets at cooler T_{eff} values do not affect our arguments concerning *turnoff* star colors and ages.

cerning the accuracy of isochrones in the region of low mass stars.

However, we do make note of some observations concerning the lower main sequences of the Yale and Victoria isochrones. We emphasize again that, because of the uncertainties in this region of the H-R diagram, we are not making an attempt to evaluate the accuracy of the Victoria or the Yale isochrones at these low temperatures. We merely present these points for completeness:

(1) First, if we take the subdwarf Gmb 1830 and adopt a color $(B - V) = 0.75$ (as V92 does), and assume $[\text{Fe}/\text{H}] = -1.30$ (as V92 does and consistent with the results of Smith *et al.* 1992), and take $\log g = 4.53$ (from Tomkin *et al.* 1992), the derived T_{eff} from the Victoria transformations (as given in Table 3 of V92) is 5087 K. The Yale T_{eff} relations yield 5130 K if one uses the transformations given by Table 1 in G88 and assumes the Vandenberg & Bell (1985) $(B - V)$ colors have been reddened by 0.025 mag to form the transformations in Table 1 of V92. The Yale color conversion result agrees slightly better with the value from the relations in chapter 2.

(2) The data points in Figure 13 show the M92 fiducial from Stetson & Harris (1988). The solid line is the 14 Gyr, $[\text{Fe}/\text{H}] = -2.26$, $Y = 0.235$ O-enhanced isochrone from BV92. As in Figure 15 of BV92, the isochrone has been adjusted vertically such that $(m - M)_V = 14.78$. The isochrone was adjusted horizontally, 0.049 mag redward, so that the observed and theoretical sequences were aligned at ~ 0.5 mag below the turnoff. As one proceeds down the main sequence, the isochrones and fiducial become increasingly divergent (though the offset is rather small). This might be expected if the redward shift applied to the Vandenberg & Bell (1985) colors by V92 was not large enough as implied by the residuals in Table 1 of G88 and the temperature scale of chapter 2. If true, then one probably expects better agreement between the Revised Yale Isochrones and the data since these colors are presumably redder at lower T_{eff} .

Figure 14 shows the same M92 fiducial and the 16 Gyr, O-enhanced, $[\text{Fe}/\text{H}] = -2.29$, $Y = 0.24$ Yale isochrone which makes use of the new G88 color transformations. This isochrone is also shown in Figure 5 of Chaboyer *et al.* (1992). The vertical shift of the isochrone follows Figure 5 of Chaboyer *et al.* in that $(m - M)_V = 14.65$. A slightly larger redward horizontal shift, 0.034 mag, than that used by Chaboyer *et al.* was made to match the isochrone to the fiducial as described above. In the range $0.55 \leq (B - V) \leq 0.85$, the Yale isochrone is a better match to the lower M92 main-sequence. Again, though, any conclusions regarding the accuracy of the RYI and the Victoria isochrones based on this comparison would be premature. Below $(B - V) \sim 0.8$, the new Yale isochrones seem to diverge from the observed fiducial and the Victoria isochrone is in better agreement with the observations. We do not know the cause of this discrepancy, but the difference is not consistent with the color

transformations being the cause. We speculate that this could be due to systematic effects introduced by the photometric analysis of these faint objects.

7. Conclusions and Summary

We have investigated possible cosmological implications of the hotter T_{eff} scale for halo stars suggested in chapter 2. The spectroscopic T_{eff} implied for the halo star HD 89499 using both experimental and theoretical $\log gf$ values is in excellent agreement with the photometric estimate from the revised temperature scale. Our T_{eff} values are also found to be consistent with those derived from the Bell & Gustafsson (1989) relations and with the T_{eff} values recently derived for two stars from Balmer line fitting. Surprisingly, the temperatures for near-turnoff and subgiant stars estimated from the color- T_{eff} transformations of both the Victoria group's and the Yale group's isochrones are in very good agreement with the scale from chapter 2. Because these transformations were derived independently, we take this agreement as good evidence for the accuracy of the previous chapter's T_{eff} scale. The next generation of large telescopes will enable abundances to be obtained in globular cluster turnoff stars. If these are to provide more reliable abundances in order to better constrain absolute or relative ages of globular clusters, attention will have to be paid to consistency between the T_{eff} scale of theoretical isochrones and that used in the derivation of abundances.

The new Yale color- T_{eff} transformations are redder than the Victoria transformations for cooler main sequence stars. Comparison of the isochrones with the observed main sequence of M92 indicates that the Yale transformations are in better agreement with the observations down to $(B - V) \sim 0.8$. This agreement with the observations is consistent with the finding in chapter 2 that cooler ($4900 \leq T_{\text{eff}} \leq 5500$ K) halo stars seemed to require the largest upward revisions of their temperatures. However, given the uncertainties of the lower main sequence models, this is a rather tentative conclusion and really does not speak to the issue of which set of isochrones is more accurate for lower mass stars. Since the red giant branches of both the Yale and the Victoria isochrones are forced to match the observed sequences of Frogel, Persson & Cohen (1981) (whose T_{eff} scale is consistent with that suggested here), the isochrones' T_{eff} values for red giants are also consistent with the values from the scale of chapter 2. We conclude that no revision in globular cluster turnoff ages is necessary for those clusters having $[\text{Fe}/\text{H}] \sim -1.3$. We can not draw the same conclusion for clusters having other metallicities. However, the good agreement between the T_{eff} values derived from the relations in chapter 2 and the Victoria transformations suggest that revision in ages of clusters of other metallicities is unlikely.

The higher T_{eff} scale suggested here leads to a halo star ${}^7\text{Li}$ plateau abundance of $N(\text{Li}) = 2.28$ from which a conservative estimate of the primordial abundance, $N(\text{Li})_{\text{p}} = 2.35$, is made. This ${}^7\text{Li}$ abundance is not simultaneously consistent with both the primordial ${}^4\text{He}$ abundance itself given by Fuller *et al.* (1991) and Boesgaard & Steigman's (1985) $\text{D}+{}^3\text{He}$ abundance limit within the framework of the standard model of Big Bang nucleosynthesis. However, if one includes both the generous observational and model uncertainties, then consistency can be achieved in the range $0.012 \leq \Omega_{\text{b}} \leq 0.017$ (for $H_0 = 100$ km/s/Mpc). The 2σ upper limit to the ${}^7\text{Li}$ abundance yields $\Omega_{\text{b}} \leq 0.019$ (or $\Omega_{\text{b}} \leq 0.12$ for $H_0 = 40$ km/s/Mpc) and so previous conclusions about the need for substantial amounts of non-baryonic matter in order to maintain consistency with inflationary models are not affected.

The T_{eff} value for HD 140283 suggested here, 5813 K, results in a B/Be ratio for this star of 12. This ratio is slightly higher than the value of 10 determined by Duncan *et al.* (1992), who adopted a lower temperature. The ratio of 12 is more comfortably consistent with the predictions of spallation production models, which predict a B/Be value of ≥ 10 . Hence, the T_{eff} scale of chapter 2 lessens the chance that the B and Be observed in HD 140283 has a significant primordial component. However, only future observations of the B/Be ratio in more metal-poor stars will be able to rule out significant primordial abundances of these elements predicted by some non-standard models of Big Bang nucleosynthesis which contain inhomogeneities. We noted that the B/Be ratio for HD 140283 predicted by the Duncan *et al.* (1992) model assumed $[\text{O}/\text{Fe}] = +0.5$ for this star. Such a low $[\text{O}/\text{Fe}]$ value has never been obtained for this mild subgiant or subgiants and dwarfs of similar metallicity. However, if one adopts the T_{eff} scale suggested here, then $[\text{O}/\text{Fe}] = +0.50$ results. Fortunately, as Duncan *et al.* (1992) point out and as is confirmed here, the predicted B/Be ratio is very insensitive to the assumed $[\text{O}/\text{Fe}]$ ratio. Indeed, changing the assumed (O/Fe) ratio by a factor of three resulted in only an $\sim 5\%$ change in the predicted B/Be ratio.

Table 7. Rederived Halo Star ${}^7\text{Li}$ Abundances

Star	[Fe/H]	T_{eff} (K)	$N(\text{Li})$
HD 16031	-1.78	6180	2.28
HD 19445	-2.06	6010	2.27
HD 74000	-2.02	6220	2.16
HD 84937	-2.21	6315	2.18
HD 134169	-0.78	5867	2.34
HD 140283	-2.57	5813	2.32
HD 201891	-1.00	6013	2.24
HD 219617	-1.36	6047	2.39
BD-10 388	-2.39	6161	2.34
BD+2 3375	-2.35	5990	2.17
BD+3 740	-2.76	6255	2.17
BD+9 352	-1.84	6007	2.26
BD+24 1676	-2.36	6341	2.36
BD+26 2606	-2.61	6114	2.31
BD+29 2091	-1.91	5820	2.25
BD+34 2476	-2.08	6280	2.25
BD+42 2667	-1.34	6074	2.23
BD+72 94	-1.61	6414	2.47
G48-29	-2.66	6420	2.34
G59-24	-2.44	6121	2.34
G64-12	-3.23	6310	2.25

Table 8. Halo Star T_{eff} Comparison

Star	[Fe/H]	$\log g$	T_{eff} K93	T_{eff} V92	ΔT_{eff} V92–K93	T_{eff} G88	ΔT_{eff} G88–K93
HD 16031	-1.78	3.92	6180	6185	+5	6086	−94
HD 19445	-2.06	4.40	6007	6018	+11	6015	+8
HD 64090	-1.66	4.75	5523	5511	−12	5551	+32
HD 74000	-2.02	4.35	6220	6222	+2	6218	−2
HD 188510	-1.52	5.05	5613	5621	+8	5686	+73
BD+29 2091	-1.91	4.33	5820	5840	+20	5863	+43
BD+34 2476	-2.08	3.67	6280	6259	−21	6263	−17
BD+72 94	-1.61	4.22	6414	6413	−1	6390	−24

Figure 11. a) Iron abundances derived by Balachandran *et al.* (1993) for each of their Fe I lines using solar gf values plotted against the excitation energy of the transition corresponding to each line. The dashed line denotes the mean abundance. Note the lack of any significant correlation. b) The same as (a) except the iron abundances have been adjusted to reflect the values that would have been found by Balachandran *et al.* (1993) using the theoretical gf values of Kurucz. There is a significant correlation between the $[\text{Fe}/\text{H}]$ values and χ .

Figure 12. a) The $\pm 2\sigma$ range for the primordial ^4He abundance (short-dash curves) predicted by the calculations of Krauss & Romanelli (1990) is plotted against the logarithm of η (multiplied by 10^9), the nucleon-to-photon ratio. The solid vertical line to the left indicates the value of η implied by a primordial ^4He abundance of $Y = 0.22$ (Fuller *et al.* 1991). The solid vertical line on the right indicates the minimum value of η implied by the $\text{D}+^3\text{He}$ abundance limit of Boesgaard & Steigman (1985). b) The solid curve shows the predicted primordial ^7Li abundance, $\log N(\text{Li})/N(\text{H})$, from the calculations of Krauss & Romanelli (1990). The $\pm 2\sigma$ range of uncertainty in these calculations is denoted by the short-dashed curves. The horizontal, long-dashed line indicates the primordial ^7Li abundance found in this paper. The two vertical solid lines have the same meaning as in (a).

Figure 13. The M92 fiducial sequence (filled squares) from Stetson & Harris (1988) is compared with a O-enhanced isochrone (solid line) from Bergbusch & Vandenberg (1992). The isochrone is parameterized by the values shown in the upper right of the panel. The isochrone has been shifted vertically and horizontally by the amounts shown in the panel.

Figure 14. The same as figure 13 except the M92 fiducial sequence is compared with the O-enhanced Yale isochrone (solid line) from Chaboyer *et al.* (1992).

Chapter 4. The Location of the Break in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ Plane

Abstract

We examine the question of the location of the break in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ relation. Fitting simple functions to a relatively homogeneous data set, we find that goodness of fit does not currently favor any unique break point in the range $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$. Statistical analysis of fits to existing ^9Be data, which we use as a proxy for O, results in the same conclusion. However, we find other evidence which does suggest that the break occurs near $[\text{Fe}/\text{H}] \sim -1.0$. This is higher than recently claimed values near $[\text{Fe}/\text{H}] \sim -1.7$. However, breaks at intermediate $[\text{Fe}/\text{H}]$ values cannot be excluded. Some of the qualitative implications of this result are discussed with particular attention to the possibility of significant delays between the end of halo formation and the beginning of star formation in the disk. A gap in the observed $[\text{O}/\text{H}]$ distribution may betray the occurrence of such a hiatus. The gap suggested by other workers to be present seems to persist despite the addition of more O data. However, the observed gap seems to be related to the well-known shortfall of stars with $-1.2 \leq [\text{Fe}/\text{H}] \leq -0.8$. More work is still required to discern whether this latter shortfall is only a selection effect or related to a genuine “G dwarf problem”. We argue that present data indicate that $[\text{O}/\text{Fe}]$ is constant in the halo (at least for $[\text{Fe}/\text{H}] \gtrsim -3$). For $[\text{Fe}/\text{H}] \geq -1.0$, ordinary least squares regressions indicate that the slope in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ relation is -0.35 to -0.40 , a fair amount less than the value (~ -0.5) usually assumed. However, the ordinary least squares bisector regression, which may be more appropriate for determining the relation between $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$, yields a steeper slope (-0.7). Until a consensus can be reached on which regression is applicable for what purposes, the constraints provided by an “observed” $[\text{O}/\text{H}]$ –age relation (which does not use true O abundances, but adopts O abundances based on Fe abundances) on several issues of Galactic chemical evolution should be regarded with caution.

1. Introduction

Spectroscopic abundance analyses of the past 15 years have revealed that stellar $[\text{O}/\text{Fe}]$ ratios rise from $[\text{O}/\text{Fe}] \sim 0$ at $[\text{Fe}/\text{H}] = 0$ to larger values (the precise value is the subject of some debate, but is near $+0.5$) in halo stars. In these more metal-poor stars, the $[\text{O}/\text{Fe}]$ ratio seems to level off to a roughly constant value (though this constancy has also been questioned). A commonly held belief concerning stellar O abundances and their import for Galactic chemical evolution is that the break in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ relation occurs very close to $[\text{Fe}/\text{H}] = -1.0$. This inference has been

supported by results for other α -elements such as Mg and S (see Wheeler, Sneden & Truran 1989 for a good review of elemental abundance trends). The notion of a break in the trend of $[O/Fe]$ vs. $[Fe/H]$ occurring at $[Fe/H] = -1.0$ has, however, recently been challenged. Bessell, Sutherland & Ruan (1991, hereafter BSR) determined O abundances in several metal-poor dwarfs from near-UV OH lines. They suggest that the break point in the $[O/Fe]$ vs. $[Fe/H]$ relation occurs at $[Fe/H] \leq -1.7$ and not at $[Fe/H] = -1.0$. Matteucci & Francois (1992, hereafter MF) argue that a break occurring at $[Fe/H] = -1.7$ can be reproduced by a detailed model of Galactic chemical evolution.

Determining the location of the break in the $[O/Fe]$ vs. $[Fe/H]$ relation has important implications for Galactic formation and chemical evolution. The general behavior of $[O/Fe]$ vs. $[Fe/H]$ is commonly explained as follows: O is synthesized and dispersed by Type II supernovae whereas iron production occurs in Type Ia supernovae. Because Type II supernova progenitors are massive, short-lived stars while Type Ia supernovae have lower mass, longer-lived progenitors, there is a lag in the accumulation of Fe compared to O in the interstellar gas. The break in the $[O/Fe]$ vs. $[Fe/H]$ relation occurs when Type Ia supernovae begin to contribute significant amounts of Fe to star-forming interstellar gas (we also point the reader to the interesting alternative explanation of Edmunds *et al.* 1991). As pointed out by Wheeler *et al.* (1989), $[Fe/H] \sim -1$ is also commonly assumed to represent the minimum metallicity of the Galactic disk⁶. This leads Wheeler *et al.* (1989) to pose an interesting question—namely, is this a mere coincidence or does it imply that the time of formation of the disk is somehow related to the evolutionary timescale of Type Ia supernova progenitors in the halo? Smecker-Hane & Wyse (1992) note that knowledge of the location of the break in the $[O/Fe]$ vs. $[Fe/H]$ relation and the use of supernova models can provide a measure of the length of time star formation occurred in the halo. This length of time is particularly important as it would seem to constrain alleged age spreads among globular clusters⁷. Smecker-Hane & Wyse (1992) discuss other important implications (including the possibility of a significant delay between the end of star formation in the halo and the beginning of star formation in the disk) of the location of the break in the $[O/Fe]$ vs. $[Fe/H]$ relation. Additional discussion concerning the importance of locating the break point in $[O/Fe]$ can be found in Smecker-Hane & Wyse (1992) and Wheeler *et al.* (1989). However, it is worth pointing out that if the break occurs at $[Fe/H] = -1.7$, then the discrepancy between suggested age spreads in the halo and the evolutionary time scale for Type Ia supernova progenitors becomes even larger.

⁶ We neither advocate nor refute that this is the minimum metallicity of the disk or that one value is applicable to the disk as a whole.

⁷ Again, we do not necessarily advocate such age spreads here.

Unfortunately, there does not seem to be a consensus on the location of the break in $[\text{O}/\text{Fe}]$. We noted above that it has been assumed in the past that this break occurs at $[\text{Fe}/\text{H}] = -1.0$, but that BSR and MF suggest $[\text{Fe}/\text{H}] = -1.7$ as a more appropriate value. To add to the confusion, Smecker-Hane & Wyse (1992) do reference the BSR work, but claim that a break in $[\text{O}/\text{Fe}]$ at $[\text{Fe}/\text{H}] = -1.0$ is “well defined”. This paper seeks to address the issue of the $[\text{O}/\text{Fe}]$ break’s location with an analysis improving on those above on three grounds: a) Using a larger number of data points than the 9 relied upon by BSR in their figure 3 to make an assessment of the location of the $[\text{O}/\text{Fe}]$ break. b) We make use of other researchers’ data that has been re-examined and reanalyzed by BSR. The data used by MF to compare with their chemical evolution models are taken, in part, from the work of Barbuy (1988) and Barbuy & Erderlyi-Mendes (1989). BSR were “unable to reproduce the O abundances of ... [Barbuy and Barbuy & Erderlyi-Mendes] exactly”. BSR rederive these abundances, providing a more homogeneous data set. c) We attempt to use a more rigorous quantitative analysis rather than qualitative inspection. MF provide no quantitative estimates of the goodness of fit of their model to the data. Therefore, it is difficult to say what the effect of the rederived abundances would be on their conclusions and whether or not the data could be fit equally well by potential models having a break in $[\text{O}/\text{Fe}]$ at $[\text{Fe}/\text{H}] = -1.0$.

2. Analysis

2.1 The Data

In order to improve on previous determinations of the $[\text{O}/\text{Fe}]$ break’s location, we seek a large number of relatively homogeneous data. The data discussed by BSR are well-suited for this purpose. We made use of BSR’s reanalyzed Barbuy (1988) and Barbuy & Erdelyi-Mendes (1989) data which were kindly provided by Dr. M. Bessell. These data were combined with the O data of Kjaergaard *et al.* (1982) (whose results were reproduced by BSR) and the O abundance determinations, based on near-UV OH lines, of BSR themselves. The abundances from BSR’s reanalysis are homogeneous in the sense that they are derived by the same investigators, using the same software, and utilizing the same model atmospheres. The data remain inhomogeneous to the extent that they were obtained using different instrumentation, reduced by different investigators, and analyzed using stellar parameters which may have systematic study-to-study differences. It seems preferable, however, to use the rederived abundances rather than the original values in order to exploit the advantages of increased homogeneity.

Additionally, we also draw on the $[\text{O}/\text{Fe}]$ data in the reanalysis by King (1993); his results

merge nicely with the data (used here) shown in figure 2 of BSR and increase the number of data points available for study. Much of the analysis was repeated after excluding King’s (1993) results and no significant differences were found. The combined data set is shown in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ diagram of Figure 15. We note in advance of quantitative results that any inhomogeneity in our sample is not at all apparent. The 1σ dispersion in our $[\text{O}/\text{Fe}]$ ratios for $[\text{Fe}/\text{H}] \lesssim -1.0$ is ~ 0.15 dex. This is comparable to (or perhaps even slightly *less* than) the dispersion expected from errors in the abundance analyses. This provides some confidence that the use of different data sets does not produce spurious effects which might hide a real break point or create an artificial one in the combined data set.

A potential concern is the fact that, because most of the stars in the sample are giants, the O abundances could be affected by $\text{O} \rightarrow \text{N}$ cycling. However, Laird (1985) finds that only a small fraction ($\sim 10\%$) of halo stars exhibit N overabundances and that these are attributable to primordial variations. This explanation is consistent with the Li detections in three N-enriched halo dwarfs by Spite & Spite (1986). N abundances could be found in the literature for only 6 of the stars used here. The stars with relatively low O abundances— $[\text{O}/\text{Fe}] \leq 0.38$ say— do not have higher $[\text{N}/\text{H}]$ ratios than other stars with average or higher $[\text{O}/\text{Fe}]$ ratios. Nor is there any correlation between the size of any potential O depletion (as measured from the mean $[\text{O}/\text{Fe}]$ ratio at a given $[\text{Fe}/\text{H}]$) and $[\text{N}/\text{H}]$. While $\text{O} \rightarrow \text{N}$ cycling seems to be important for interpretation of O abundances in some globular cluster giants (*e.g.*, Sneden *et al.* 1992), it is apparently not important for the sample at hand. In any case, we emphasize that the objective of the present work is not to form a sample free from the potential biases of mixing processes; rather, we seek to analyze the same data which has been used before to argue for a given break point, but in a more quantitative fashion.

2.2 Methodology

We now attempt to quantitatively constrain the location of the break in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ relation. There exists a large statistical literature on methods to locate such “change-points” in trends of data. While sophisticated numerical and geometric techniques exist to address such problems, we formulate a very simple approach here. The more elaborate techniques are based on both specific assumptions concerning the distribution of errors in the data (*e.g.*, normally distributed errors) and the more general “non-parametric” assumption. Here we will assume normally distributed errors in the O and Fe abundances as we have no reason to believe otherwise. Fortunately, even those sophisticated techniques (some similar in nature to the approach used here) which make the

assumption of normally distributed errors tend to be forgiving of non-normally distributed errors.

The first method used here is as follows. We fit the $[\text{O}/\text{Fe}]$ *vs.* $[\text{Fe}/\text{H}]$ data with a simple function in two separate regimes of $[\text{Fe}/\text{H}]$. The value of $[\text{Fe}/\text{H}]$ separating these regimes is considered as the break in the $[\text{O}/\text{Fe}]$ *vs.* $[\text{Fe}/\text{H}]$ relation. One can then measure the goodness of fit by computing the value of χ^2 or, more specifically, the reduced χ^2 , $\chi^2_\nu = \chi^2/\nu$ where ν is the number of degrees of freedom. We can then compare values of χ^2 from fits assuming various break points using a standard F-test (*e.g.*, Bevington, 1969) to determine if the goodness of fit is significantly better for one assumed breakpoint than another. While this simple approach is not the most robust way of isolating the break point in $[\text{O}/\text{Fe}]$, it can be used to test the contention of MF—namely that goodness of fit (measured quantitatively or otherwise) can constrain the break point.

We must first decide what sort of function to fit to the data shown in Figure 15. Such a choice must be a physically plausible one. In this respect, the results of Galactic chemical evolution models can provide useful guidance. The models of MF indicate that for the more metal-rich stars, a roughly linear relation exists between $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$. For the more metal-poor stars (or those with $[\text{Fe}/\text{H}] \gtrsim -3$ anyway), MF's results indicate that a linear relation with a small slope would be an appropriate function to fit to our data. Thus, we will fit two lines to the two regimes in $[\text{Fe}/\text{H}]$ separated by an assumed break point.

Before proceeding, we further consider the choice of the function to be fit to $[\text{O}/\text{Fe}]$ at low values of $[\text{Fe}/\text{H}]$. The model of MF has a small slope at low $[\text{Fe}/\text{H}]$ becoming much steeper below $[\text{Fe}/\text{H}] \sim -3$. As will be discussed later, the post-primordial Galactic ^9Be abundance is dependent on the O abundance. Ryan *et al.* (1992) are able to reproduce the observed run of Be abundance *vs.* $[\text{Fe}/\text{H}]$ using a simple model of Galactic chemical evolution which assumes a *constant* $[\text{O}/\text{Fe}]$ ratio from $[\text{Fe}/\text{H}] \sim -2.7$ to $[\text{Fe}/\text{H}] = -1.0$. Using guidance provided by the O data themselves, we compute the ordinary correlation coefficient, r , using the $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$ data. We find that r has values of -0.205 , -0.173 , -0.304 for the data having $[\text{Fe}/\text{H}]$ less than -1.0 , -1.3 , and -1.7 respectively. These values indicate correlations at only the 81%, 68%, and 87% confidence levels respectively. No significant correlation is found in the full sample of $[\text{O}/\text{Fe}]$ ratios presented by Tomkin *et al.* (1992) or in a subsample of theirs reanalyzed by King (1993). While the model of MF may fit existing O data just as well as alternative models having a constant $[\text{O}/\text{Fe}]$ ratio in the halo, the data themselves show no statistically significant trend between $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$. Because O data in the extreme halo ($[\text{Fe}/\text{H}] \leq -3$) will be able to provide further and better constraints on Galactic chemical evolution calculations, such models should presently be regarded with caution.

It should be noted that we are fitting $[\text{O}/\text{Fe}]$, not $[\text{O}/\text{H}]$, *vs.* $[\text{Fe}/\text{H}]$. This raises the concern of

potential correlation of errors in the O and Fe abundances which could affect the analysis. There seems to be no reason to believe that the errors in the O and Fe abundances would be correlated for the stars drawn from BSR. The same statement might not be true for the stars drawn from the reanalysis of King (1993). In this case, errors in the O and Fe abundances due to uncertainties in the adopted T_{eff} would be correlated. However, the stars from King (1993) constitute a small portion of our entire sample. In order to directly address this concern, much of the analysis was repeated in the $[\text{O}/\text{H}]$ vs. $[\text{Fe}/\text{H}]$ plane with essentially identical results. Therefore, we will continue to work in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ plane here as this is the one commonly used by and familiar to most astronomers.

2.3 Linear Regressions

2.3.1 Which Regression to Use?

Most linear regressions performed in astronomical applications are of the ordinary least squares (OLS) variety where the sum of the square of the residuals in the dependent variable (in this case $[\text{O}/\text{Fe}]$) are minimized. In two recent papers, Isobe *et al.* (1990, hereafter IFAB) and Feigelson & Babu (1992, hereafter FB) caution against indiscriminate use of this method. As they emphasize, careful consideration and a precise statement of the problem at hand should dictate which form of regression is most suitable.

In the remainder of this section we are interested in estimating, using the wording of IFAB, “the underlying functional relation between” $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$. In this case, IFAB recommend using the OLS bisector method which has the smallest variance about theoretical fits among methods which treat the independent and dependent variables symmetrically. We use the OLS bisector regression for the data in the high $[\text{Fe}/\text{H}]$ regime. Unfortunately, the OLS bisector does not give very realistic results for data in the low $[\text{Fe}/\text{H}]$ regime. This can be understood as follows. The OLS bisector regression yields the line which bisects the $\text{OLS}([\text{O}/\text{Fe}][\text{Fe}/\text{H}])$ and $\text{OLS}([\text{Fe}/\text{H}][\text{O}/\text{Fe}])$ lines. Because of the nature of the latter regression, the best fitting line will be a precisely *vertical* line in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ plane if $[\text{O}/\text{Fe}]$ is precisely constant at low $[\text{Fe}/\text{H}]$. The resulting OLS bisector regression of $[\text{O}/\text{Fe}]$ data that is precisely constant at low $[\text{Fe}/\text{H}]$ will then be a line having a 45° degree slope. Since our $[\text{O}/\text{Fe}]$ data are roughly constant at low $[\text{Fe}/\text{H}]$, such an implausible regression given by the OLS bisector method in this special case is obtained. In order to avoid this dilemma we must choose another regression that treats the $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$ data symmetrically. The reduced major axis regression described by IFAB gives no information on the intrinsic relation between $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$. Thus, for fitting data in the low $[\text{Fe}/\text{H}]$ regimes, we

choose to use the orthogonal regression method, described by IFAB, which minimizes the residuals measured perpendicular (orthogonally) to the regression line.

In §4 we will be interested in predicting an $[\text{O}/\text{Fe}]$ value at a given $[\text{Fe}/\text{H}]$ value using a regression calculated for data in a different $[\text{Fe}/\text{H}]$ regime. As IFAB note, this specific problem favors use of an $\text{OLS}([\text{O}/\text{Fe}][\text{Fe}/\text{H}])$ regression. The discussion in IFAB indicates that an OLS regression of $[\text{O}/\text{Fe}]$ on $[\text{Fe}/\text{H}]$ – $\text{OLS}([\text{O}/\text{Fe}][\text{Fe}/\text{H}])$ – assumes that: a) the errors in $[\text{Fe}/\text{H}]$ are zero b) the errors in $[\text{O}/\text{Fe}]$ do not depend on $[\text{Fe}/\text{H}]$ and c) the errors $[\text{O}/\text{Fe}]$ have a common, independent variance. Certainly, assumption a) does not hold. It would also seem that b) and c) are doubtful. As one goes to more metal-poor stars, the S/N of the spectra are probably lower (leading to uncertainties in measuring the line strength and setting the continuum), the O line strengths are weaker, *etc.*, making the O abundances more uncertain. Despite the possible violations of the above assumptions, IFAB suggest that the OLS method is still valid as long as the regression will be used to predict $[\text{O}/\text{Fe}]$ given a value of $[\text{Fe}/\text{H}]$, which is the case in §4.

It turns out that identical conclusions are reached in §2.4 whether one uses OLS or OLS bisector/orthogonal regressions. We will present the results using both regressions as the OLS fits and χ^2_ν values could probably be more easily verified and reproduced by other workers. We caution, though, that we cannot justify the use of OLS regressions *per se* in the analysis of §2.4. Dr. E. Feigelson kindly provided the software, SLOPES, used to perform the OLS bisector and orthogonal regressions and calculate the respective uncertainties. The $\text{OLS}([\text{O}/\text{Fe}][\text{Fe}/\text{H}])$ regressions were performed using the IRAF package and private software. However, the OLS slopes and intercepts determined by the SLOPES package turn out to be identical. This is not the case for the uncertainties in these quantities. Our OLS uncertainties were calculated based on the methods in, *e.g.*, Bevington (1969). FB point out that the variances calculated by this method are likely to be mathematically incorrect. Indeed, the OLS variances determined by the SLOPES software are 10 – 15% lower than those determined using the methods in Bevington (1969). However, FB note that for small samples ($N \leq 50$ say, which is the case here), SLOPES' variances are likely to be underestimates of the true uncertainties in the fits. While OLS variances determined by the method of Bevington (1969) may be formally incorrect, they are probably more realistic and we retain them for the OLS fits.

2.3.2 Weighted and Unweighted Fits

The regressions described above are unweighted regressions. These are appropriate if only information on $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$ and not their uncertainties is available or if the intrinsic scatter in

[O/Fe] and [Fe/H] is unknown. The latter situation could be argued to hold *a priori*. If halo collapse was not a single rapid event, but instead was a delayed accumulation of independent fragments (each with a distinct history of chemical evolution), then appreciable scatter might be expected in [O/Fe] ratios at low metallicity. Similarly, if the initial mass function or star formation rate varies as a function of radius, azimuth, or height in the disk, then no tight relation between [O/Fe] and [Fe/H] may exist at higher metallicities. In these cases, unweighted structural regression is appropriate.

However, information on the measurement uncertainties is available. The scatter about the mean [O/Fe] for stars having $[\text{Fe}/\text{H}] \leq -1.0$ is ~ 0.16 dex, which is consistent with the measurement uncertainties (~ 0.20 dex). In this case, weighted functional regression is appropriate. As discussed in FB, such regressions fall into two broad categories— homoscedastic and heteroscedastic regression models. The former assumes identical point-to-point scatter resulting from measurement error. The latter assumes different point-to-point measurement errors.

Although heteroscedastic regressions are desirable for the problem at hand, we are forced to use a homoscedastic regression. This is due to the lack of knowledge of the precise errors in [O/Fe] and [Fe/H] for each individual data point. Most of the analyses from which our data are drawn make no estimate of the uncertainty in [Fe/H] and generally assign a single uncertainty to all the [O/Fe] ratios. The homoscedastic regressions make use of the parameter λ , defined here to be $\sigma^2([\text{O}/\text{Fe}])/\sigma^2([\text{Fe}/\text{H}])$. Values of $\sigma = 0.15$ dex were assumed for both [O/H] and [Fe/H] ratios. This results in $\sigma([\text{O}/\text{Fe}]) = 0.21$ and so $\lambda = 2$. Least squares fits were carried out using software based on equation (1) in FB. Identical conclusions (though of course, not identical results— the exact values of the individual reduced χ^2 are different) are reached using these fits. In the interest of brevity and because readers will probably be most familiar with the simple unweighted regressions and have access to software to verify these results, we present only the results of the unweighted fits.

3. Results

Fitting a line of the form $[\text{O}/\text{Fe}] = a[\text{Fe}/\text{H}] + b$ (where a and b are constants determined in the fit) for the data with $[\text{Fe}/\text{H}] \geq -1.0$ yields⁸ $\chi^2_\nu = 1.245$ (OLS fit, 0.903). A similar fit to the data with $[\text{Fe}/\text{H}] \leq -1.0$ yields $\chi^2_\nu = 0.579$ (OLS fit, 0.579 also). The combined reduced χ^2 for the whole fit is 0.941 (OLS, 0.755). Forcing [O/Fe] to be a constant (taken as the straight mean of the values— in this case $[\text{O}/\text{Fe}] = 0.462$) for $[\text{Fe}/\text{H}] \leq -1.0$ gives a combined χ^2_ν of 0.942 (OLS 0.758). The combined fit χ^2 values for the two assumptions about the behavior of [O/Fe] at low metallicities also indicate

⁸ In the actual computation of χ^2_ν we assume the uncertainty in [O/Fe] to be ± 0.20 dex.

that there is no reason to favor increasing $[\text{O}/\text{Fe}]$ values with decreasing $[\text{Fe}/\text{H}]$ at low $[\text{Fe}/\text{H}]$.

We find that $\chi^2_\nu = 0.925$ (OLS, 0.796) for the fit of the data having $[\text{Fe}/\text{H}] \geq -1.7$. The corresponding value is $\chi^2_\nu = 0.631$ (OLS, 0.628) for the data having $[\text{Fe}/\text{H}] \leq -1.7$. The combined reduced χ^2 for the whole fit is $\chi^2_\nu = 0.845$ (OLS, 0.750). Taking a constant $[\text{O}/\text{Fe}]$ ratio of 0.464 (determined again by a straight mean) for data having $[\text{Fe}/\text{H}] \leq -1.7$ results in $\chi^2_\nu = 0.852$ (OLS, 0.759) for the combined fit over all $[\text{Fe}/\text{H}]$.

A fit to the data with $[\text{Fe}/\text{H}] \geq -1.3$ yields $\chi^2_\nu = 1.125$ (OLS, 0.881) while a fit to our data having $[\text{Fe}/\text{H}] \leq -1.3$ yields $\chi^2_\nu = 0.579$ (OLS, 0.578). The combined reduced χ^2 for both fits is $\chi^2_\nu = 0.906$ (OLS, 0.760). Using the mean value of $[\text{O}/\text{Fe}] = 0.470$ for the data with $[\text{Fe}/\text{H}] \leq -1.3$ gives a combined reduced χ^2 for the fits in both $[\text{Fe}/\text{H}]$ regimes of $\chi^2_\nu = 0.903$ (OLS, 0.759).

We can now compare the above values of χ^2_ν for the combined fits using a standard F-test to determine if any of the differences are significant. Such an analysis of the goodness of the fits (including several, not detailed here, which assumed various break points between $[\text{Fe}/\text{H}] = -1.7$ and -1.0) reveals that none are statistically better than another; the F tests all indicate significant differences between the various fits at only the $< 75\%$ confidence level. We emphasize that *goodness of fit to current $[\text{O}/\text{Fe}]$ data does not favor any unique break point over another in the range $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$.*

A finding of some interest concerns the slopes for the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ relation at high $[\text{Fe}/\text{H}]$. The values given by the OLS fits are $-0.363 (\pm 0.079)$, $-0.372 (\pm 0.060)$, and $-0.347 (\pm 0.036)$ (where the values in parentheses are the 1σ uncertainties based on the formal variance of the fitted value) for $[\text{Fe}/\text{H}] \geq -1.0$, -1.3 , and -1.7 respectively. These slopes are consistently shallower than the value usually assumed—namely, $[\text{O}/\text{Fe}] \sim -0.5[\text{Fe}/\text{H}]$ (Wheeler *et al.* 1989). However, the correct regression to use if one wants to know “what is the relation between $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$?” is the OLS bisector. The slopes yielded by this regression are $-0.703 (\pm 0.076)$ and $-0.605 (\pm 0.064)$ —steeper than the canonical value of -0.5 —for $[\text{Fe}/\text{H}] \leq -1.0$ and -1.3 .

It would be of some interest to know the slope predicted by detailed chemical evolution models such as those of MF. The OLS slopes for the relations having $[\text{Fe}/\text{H}] \leq -1.0$, -1.3 , and -1.7 are $-0.060 (\pm 0.045)$, $-0.057 (\pm 0.054)$, and $-0.132 (\pm 0.0824)$. The slopes and uncertainties from the orthogonal regressions are quite similar. Comparing the formal uncertainties in parentheses to the values themselves, it is apparent that none of the fitted slopes can be said to be significantly different from 0. This again calls into question the need for believing that $[\text{O}/\text{Fe}]$ is anything but constant in the halo (at least for $[\text{Fe}/\text{H}] \gtrsim -3$).

4. Additional Fits

A sample OLS fit ($[\text{O}/\text{Fe}][\text{Fe}/\text{H}]$) for the data in Figure 15 is shown in Figure 16. In this case, fits were made to the data having $[\text{Fe}/\text{H}] \geq -1.3$ and to the data having $[\text{Fe}/\text{H}] \leq -1.3$. Not all the fits having other break points resulted in the very nearly exact convergence of the fitted lines at the assumed break point as is the case at $[\text{Fe}/\text{H}] = -1.3$ shown in Figure 16. However, in all cases the offset of the fitted lines could not be regarded as significant (most fits had differences at the breakpoint differing significantly at only the $1 - 1.5\sigma$ level) given the uncertainties in the slopes and zero points.

We investigated this matter further by conducting fits which: force the lines to converge at the assumed break point, fit only the slope in the high $[\text{Fe}/\text{H}]$ regime, and assume a constant $[\text{O}/\text{Fe}]$ in the low $[\text{Fe}/\text{H}]$ regime. The reduced χ^2 of the combined fits assuming break points at $[\text{Fe}/\text{H}] = -1.7$, -1.3 , and -1.0 are 0.809, 0.755, and 0.772. As for the other fits discussed in §2.4, none of these χ^2_ν values is significantly better than another. Nor are any of these fits significantly better than those discussed in §2.4.

5. Constraints on the Break Point Provided by ^9Be Abundances

There is good reason to believe that the chemical evolution of the Galactic ^9Be abundance is related to the evolution of the O abundance (*e.g.*, Ryan *et al.* 1992). Non-primordial ^9Be production is thought to be due to spallation reactions involving cosmic rays and CNO nuclei. Since O is the most abundant of these three metals, we consider just the O abundance. The flux of cosmic rays in the halo should be proportional to the Type II supernova rate. Since halo stars which become Type II supernovae are also those that produce O in the halo, one would expect that the abundance of ^9Be in the halo is proportional to the square of the O abundance. If ^9Be is synthesized in close physical proximity to the Type II supernova (again, the stars which are responsible for O synthesis in the halo), then the halo ^9Be abundance is expected to be proportional to simply the O abundance. This latter case is apparently what is supported by ^9Be observations (*e.g.*, Boesgaard & King 1993, Gilmore *et al.* 1992). Despite the specific case which holds in the halo, in the logarithmic abundance plane the ^9Be abundance should be proportional to the O abundance with a slope between 1 and 2. Since the $[\text{O}/\text{Fe}]$ ratio appears to be constant in the halo at metallicities of interest to us here, $[\text{Be}/\text{H}]$ should be proportional to $[\text{Fe}/\text{H}]$ with a slope between 1 and 2. Thus, we can use the Be abundance as a proxy for the O abundance and investigate the location of possible break points in the Be-Fe plane.

Our ^9Be data set is composed of those stars analyzed by Boesgaard & King (1993), Gilmore *et al.* (1992), and Ryan *et al.* (1992). In the case of overlap, mean values were simply used. The data are plotted in Figure 17, which shows $\log n(\text{Be})/n(\text{H})$ vs. $[\text{Fe}/\text{H}]$. We use the same approach as for O (except the OLS regression results will not be presented), making separate linear fits to data both above and below a given $[\text{Fe}/\text{H}]$ value. It is clear that the (Be/H) vs. $[\text{Fe}/\text{H}]$ relation flattens off at high metallicity. As for the O data at low metallicity, this flatness leads to rather unrealistic regressions using the OLS bisector. Hence we will use orthogonal regressions for the high metallicity regime and OLS bisector regressions for the low metallicity regime. Although this heterogeneous approach is probably undesirable, we argue that it is justified (both for the O and Be data) because: a) it is consistently used for all the fits having different assumed break points and b) a homogeneous approach using orthogonal regression (which meets the goal of symmetric treatment of the variables) for both metallicity regimes yields the same results— *i.e.*, the slopes and intercepts of the orthogonal and OLS bisector regressions agree to well within the uncertainties for the low metallicity Be data and high metallicity O data.

The combined (*i.e.*, in both regimes of $[\text{Fe}/\text{H}]$) fits for assumed break points at $[\text{Fe}/\text{H}] = -1.0$, -1.3 , and -1.7 yield values of χ^2_ν of 0.776, 0.751, and 0.763 respectively (in computing the reduced χ^2 , we adopted an uncertainty in $\log n(\text{Be})/n(\text{H})$ of ± 0.25 dex). Comparison via an F test reveals that none of the combined fits are significantly better than another. Thus, goodness of fit to ^9Be data also does not favor any assumed changepoint over another in the range $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$.

6. Current Best Estimate of the Break in $[\text{O}/\text{Fe}]$

Since our efforts above were not able to define the location of the break point in $[\text{O}/\text{Fe}]$, we attempt to do so in a different manner. If the break in $[\text{O}/\text{Fe}]$ occurs at, say, $[\text{Fe}/\text{H}] = -1.7$, then a fit to the data having $[\text{Fe}/\text{H}] \geq -1.0$ should produce an equally good fit to the data having $-1.7 \lesssim [\text{Fe}/\text{H}] \lesssim -1.0$. We also can use the ^9Be data to perform the same analysis. We begin with the ^9Be data and utilize OLS fits for reasons outlined in §2.3.1.

Fitting the ^9Be data having $[\text{Fe}/\text{H}] > -1.0$ shown in Figure 17 with a straight line results in a reduced χ^2 of 0.722. Applying this same fit to the ^9Be data having $-1.75 \leq [\text{Fe}/\text{H}] \leq -1.0$ yields $\chi^2_\nu = 2.107$. The difference in goodness of fit is significant at the $\sim 98.5\%$ confidence level. Alternatively, fitting the ^9Be data having $[\text{Fe}/\text{H}] < -1.75$ results in $\chi^2_\nu = 0.648$. Applying this same fit to the data having $-1.75 \leq [\text{Fe}/\text{H}] \leq -1.0$ results in $\chi^2_\nu = 0.743$. This difference is significant at only $< 75\%$ confidence level. These results indicate that the data having $-1.75 \leq [\text{Fe}/\text{H}] \leq -1.0$ are

fit as equally well using the relation found for the data with $[\text{Fe}/\text{H}] < -1.75$, but are fit significantly worse by the relation found for the data with $[\text{Fe}/\text{H}] > -1.0$. This seems to suggest that the break in the Be *vs.* Fe relation occurs more closely to $[\text{Fe}/\text{H}] = -1.0$ than $[\text{Fe}/\text{H}] = -1.7$. Additional Be abundances for stars having $-2 \lesssim [\text{Fe}/\text{H}] \lesssim -1$ would be desirable to confirm this conclusion.

Turning to the O data, the fit in §2.5 which assumed a break point at $[\text{Fe}/\text{H}] = -1.0$, a constant $[\text{O}/\text{Fe}]$ at $[\text{Fe}/\text{H}] \leq -1.0$, and convergence of the two fitted lines at $[\text{Fe}/\text{H}] = -1.0$ yielded a reduced χ^2 of 0.993 for the data having $[\text{Fe}/\text{H}] \geq -1.0$. Applying this same fit for $[\text{Fe}/\text{H}] \geq -1.0$ to the data with $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$ results in a reduced χ^2 of 2.082. An F test reveals that the two χ^2_ν values are different at the $\sim 98\%$ confidence level. This significant difference suggests that the same function which can fit the data for $[\text{Fe}/\text{H}] \geq -1.0$ does not fit the O data having $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$. This is unexpected if the break point in $[\text{O}/\text{Fe}]$ lies at $[\text{Fe}/\text{H}] \sim -1.7$. Similarly, the constant $[\text{O}/\text{Fe}]$ ratio fit to the data having $[\text{Fe}/\text{H}] \leq -1.7$ yielded $\chi^2_\nu = 0.666$. Extending this constant $[\text{O}/\text{Fe}]$ ratio (0.464 ± 0.163) to the data having $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$ yields $\chi^2_\nu = 0.418$. The difference between the reduced χ^2 values (0.666 and 0.418) is not statistically significant. Again, this seems unexpected if the $[\text{O}/\text{Fe}]$ breakpoint is at $[\text{Fe}/\text{H}] = -1.7$, above which the $[\text{O}/\text{Fe}]$ ratios should not be constant.

Based on this analysis, there appears to be good evidence that the break in the $[\text{O}/\text{Fe}]$ ratio occurs at $[\text{Fe}/\text{H}] = -1.0$ rather than at $[\text{Fe}/\text{H}] = -1.7$. However, it must be emphasized that values in between cannot be excluded at this time. Additional data in all metallicity regimes, attention to corresponding kinematic data, and closer analysis of the cause of the scatter seen in Figure 15 will be needed before more firm conclusions can be drawn. The best current evidence does suggest, though, that the break in $[\text{O}/\text{Fe}]$ more likely occurs closer to $[\text{Fe}/\text{H}] = -1.0$ than $[\text{Fe}/\text{H}] = -1.7$.

7. Summary and Conclusions

This chapter examined the question of the location of the break point in the $[\text{O}/\text{Fe}]$ *vs.* $[\text{Fe}/\text{H}]$ relation. Using a relatively homogeneous data set, we fit plausible functions to the data in distinct $[\text{Fe}/\text{H}]$ regimes separated by an assumed break point. Comparing various values of the resulting χ^2_ν indicates that goodness of fit does not favor any assumed break point over another in the range $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$. Because ^9Be abundance potentially serves as a proxy for O abundance, we repeated the analysis in the $[\text{Be}/\text{H}]$ – $[\text{Fe}/\text{H}]$ plane. This comparison also indicates that goodness of fit of physically plausible functions does not favor a unique location of the $[\text{O}/\text{Fe}]$ breakpoint. Figure 18 serves to further illustrate this point. Figure 18a shows a second order polynomial least squares fit to the O data. The combined reduced χ^2 is 0.774– statistically indistinguishable from the other

OLS fits detailed in §2.4 and 2.5. Figure 18b shows a third order cubic spline fit to the data. The reduced χ^2 is 0.775, essentially identical to the second order polynomial fit. However, the break in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ fit in Figure 18a occurs near $[\text{Fe}/\text{H}] = -1.6$ whereas the break in the spline fit of Figure 18b occurs near $[\text{Fe}/\text{H}] = -1.2$.

We also find that the present data argue that $[\text{O}/\text{Fe}]$ is flat in the halo (at least for $[\text{Fe}/\text{H}] \gtrsim -3$). Fits which assume a constant $[\text{O}/\text{Fe}]$ ratio at low $[\text{Fe}/\text{H}]$ values are not significantly worse than those fits having a slope in $[\text{O}/\text{Fe}]$. Indeed, given the uncertainties in the fits, these slopes are not significantly different from zero. Furthermore, statistical analysis reveals that there is no correlation between $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$ at $[\text{Fe}/\text{H}] \leq -1.0$. Testing the predictions of MF concerning a steep increase in $[\text{O}/\text{Fe}]$ below $[\text{Fe}/\text{H}] \sim -3$ will have to await the acquisition of additional data for very metal-poor stars.

Our OLS fits to the $[\text{O}/\text{Fe}]$ data at higher metallicities suggest that the slope in $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ is between -0.35 and -0.4 . This is shallower than the commonly assumed value of -0.5 . This reduction in slope would be in better agreement with models of O nucleosynthesis in the disk such as those by Twarog & Wheeler (1982) which predict shallower slopes still at higher metallicities ($[\text{Fe}/\text{H}] \geq -0.5$). However, the OLS bisector regression is more appropriate when addressing the question “what is the relationship between $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$?”. The OLS bisector fits indicate that the slope is steeper, ~ -0.7 , than the canonical value of -0.5 . This exacerbates the discrepancy between the data used here and models for $[\text{Fe}/\text{H}] \geq -0.5$ such as those by Twarog & Wheeler (1982). These differences in the slopes (and zero points) are also important for constraining models of Galactic chemical evolution which make use of age- $[\text{O}/\text{H}]$ data (*e.g.*, Meusinger 1992a,b). Such models are almost always compared to O abundance “data” which is *not* based on observed O abundances but, rather, on transformations of the $[\text{Fe}/\text{H}]$ ratio to $[\text{O}/\text{H}]$ based on an *assumed* relation between $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$. If observed O abundances are not used in such analyses, then more attention must be paid to determining the slope and zero-point of the $[\text{O}/\text{Fe}]$ relation at high $[\text{Fe}/\text{H}]$. Specifically, it appears that astronomers will have to reach a consensus on which regression is applicable in the case of transforming $[\text{Fe}/\text{H}]$ to $[\text{O}/\text{H}]$ for the purposes of comparison to theory. Until then (or until O abundances can be obtained directly), the constraints provided by such O “data” on the issues of gas infall rate, its temporal variability, dust outflow, and dust explosion must be regarded as potentially weak.

There is some evidence that the break in the $[\text{O}/\text{Fe}]$ data is more likely to occur at $[\text{Fe}/\text{H}] = -1.0$ than at $[\text{Fe}/\text{H}] = -1.7$. We find that the fit to the $[\text{O}/\text{Fe}]$ data having $[\text{Fe}/\text{H}] \geq -1.0$ cannot fit the data having $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$. However, the constant $[\text{O}/\text{Fe}]$ value which fits the data for

$[\text{Fe}/\text{H}] \leq -1.7$ fits the data having $-1.7 \leq [\text{Fe}/\text{H}] \leq -1.0$ equally well. Analogous results are found using recent ^9Be , which we use as a proxy for O, data. These results suggest that a change in the $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ relation occurs at $[\text{Fe}/\text{H}] = -1.0$ (or at least nearer to -1.0 than -1.7). However, we emphasize that a break occurring at intermediate values cannot be currently excluded.

The implications of these results are best discussed once firmer conclusions can be reached using more data having corresponding kinematic information. However, we briefly sketch some of implications taking our conclusions at face value. *If* one wants small age spreads in the halo (for consistency with the results of some globular cluster age determinations or to match hierarchical models of halo formation via mergers of independent fragments), then the currently most plausible—in the sense of accommodating this age spread—source of Type Ia supernovae may well be CO-He white dwarf mergers (Smecker-Hane & Wyse 1992). However, as discussed by Smecker-Hane & Wyse, such a solution may not accommodate a break in $[\text{O}/\text{Fe}]$ at $[\text{Fe}/\text{H}] = -1.0$. This brings us back to the question posed by Wheeler *et al.* (1989)—why would the duration of halo formation be comparable to the evolutionary timescale of Type Ia supernova progenitors? Since no convincing arguments have been presented to suggest why this should be so, the most attractive explanation which could maintain the break at $[\text{Fe}/\text{H}] = -1.0$ is a significant delay between the end of halo formation and the beginning of star formation in the disk (Smecker-Hane & Wyse 1992).

Is there any independent evidence to suggest that there could be a hiatus between the end of star formation in the halo and the beginning of star formation in the disk? Wheeler *et al.* (1989) qualitatively demonstrate how such a hiatus can lead to gaps in the $[\text{O}/\text{H}]$ distribution. As they point out, their figure 2 suggests that such a gap may exist. In Figure 19, we show our data in the same $[\text{Fe}/\text{O}]$ vs. $[\text{O}/\text{H}]$ plane as Wheeler *et al.* (1989). It appears that despite the additional data, this gap at $-0.8 \lesssim [\text{O}/\text{H}] \lesssim -0.5$ still persists. However, determining the precise behavior of the data at higher values of $[\text{O}/\text{H}]$ will be required to discern between chemical evolution models having a hiatus and models having a burst superimposed on a constant star formation rate (which also can produce gaps in $[\text{O}/\text{H}]$).

Further examination shows that this possible gap in the $[\text{O}/\text{H}]$ distribution corresponds to the gap seen in the $[\text{Fe}/\text{H}]$ distribution of Cayrel de Strobel & Bontolila (1983). Fitting the $[\text{Fe}/\text{H}]$ vs. $[\text{O}/\text{H}]$ data used here with an OLS regression (which is preferred for the *prediction* of an $[\text{Fe}/\text{H}]$ value given a value for $[\text{O}/\text{H}]$), one finds that a gap from $-0.8 \leq [\text{O}/\text{H}] \leq -0.5$ is expected to correspond to a gap in $[\text{Fe}/\text{H}]$ in the region $-1.04 \leq [\text{Fe}/\text{H}] \leq -0.68$. Figure 20 shows the $[\text{Fe}/\text{H}]$ vs. $[\text{O}/\text{H}]$ diagram. The actual data show an apparent gap in $[\text{Fe}/\text{H}]$ in the range $-0.8 \leq [\text{Fe}/\text{H}] \leq -1.2$. This is confirmed in Figure 21, which shows the $[\text{Fe}/\text{H}]$ distribution (in bins of 0.3 dex) of the data in Figure 20. The

gap at intermediate metallicities is evident in this diagram. Tracing the smooth distribution of $[\text{Fe}/\text{H}]$ in figure 1 of Cayrel de Strobel & Bentolila (1983) using the first few metal-rich bins and the last few metal-poor bins, one finds that there is a short-fall in the bin having $-1.05 \leq [\text{Fe}/\text{H}] < -0.75$ and, to a lesser extent, the bin $-1.35 \leq [\text{Fe}/\text{H}] < -1.05$. This suggests that the gap seen in the $[\text{O}/\text{H}]$ distribution may well be associated with the long-known short-fall in the associated $[\text{Fe}/\text{H}]$ distribution. Whether this latter gap is due to a genuine “G dwarf problem” or mere selection bias is currently a matter of some debate.

Another point made by Wheeler *et al.* (1989) is worth repeating. *If* one accepts the results of Barry (1988) and Carlberg *et al.* (1985) as indicating that the age- $[\text{Fe}/\text{H}]$ relation in the disk is flat⁹, then this may indicate (but does *not require*) that there was a hiatus between the end of halo formation and the beginning of star formation in the disk. Clearly, additional O abundance data together with corresponding kinematic information will be able to address issues such as these better.

⁹ Because we do not want to become involved in the rather acrimonious debate which exists about this topic, it must be stated that these results certainly have not gone uncontested.

Figure 15. Combined data set used in the analysis plotted in the $[O/Fe]$ vs. $[Fe/H]$ plane. The abundances are taken from the data of Barbuy (1988) and Barbuy & Erdelyi-Mendes (1989) as reanalyzed by BSR; Kjaergaard *et al.* (1982); and BSR. The data from these sources are shown as filled squares. The data of Tomkin *et al.* (1992) as reanalyzed by King (1993) are shown as open squares.

Figure 16. A sample ordinary least squares fit ($[O/Fe]$ vs $[Fe/H]$) to the data shown in Figure 1. In this case, separate fits were made to the data having $[Fe/H] \geq -1.3$ (solid line) and to the data having $[Fe/H] \leq -1.3$ (dashed line). The nearly exact convergence of the two fits at the assumed break point ($[Fe/H] = -1.3$) was *not* forced.

Figure 17. Combined ^9Be data set used in the analysis plotted in the $\log n(\text{Be})/n(\text{H})$ vs. $[\text{Fe}/\text{H}]$ plane. The abundances are taken from the data of Boesgaard & King (1993), Gilmore *et al.* (1992), and Ryan *et al.* (1992).

Figure 18. a) Second order polynomial fit to the O data. The break in the fit appears to occur near $[\text{Fe}/\text{H}] = -1.6$. b) Third order cubic spline fit to the same O data. The break in this fit occurs near $[\text{Fe}/\text{H}] = -1.2$. Despite the different break points, the goodness of fit to the data of both fits is identical.

Figure 19. Data shown in Figure 15 are displayed in the $[\text{Fe}/\text{O}]$ vs. $[\text{O}/\text{H}]$ plane so that comparison can be made with figure 2 of Wheeler *et al.* (1989). The two figures taken together suggest that the gap in $[\text{O}/\text{H}]$ at $-0.8 \lesssim [\text{O}/\text{H}] \lesssim -0.5$ may well be real (in the sense of its existence being statistically significant), although its cause could be due to selection effects.

Figure 20. The data of Figure 15 are shown in the $[Fe/H]$ vs. $[O/H]$ plane. The dashed lines enclose an apparent gap, corresponding to that seen in $[O/H]$ in Figure 5, in the range of $[Fe/H]$ covered by the data. This gap in $[Fe/H]$ occurs in the same location as the shortfall in the $[Fe/H]$ distribution seen in figure 1 of Cayrel de Strobel & Bantolila (1983).

Figure 21. The $[Fe/H]$ distribution of the data shown in Figure 20. Note the pronounced dip in the number of stars at intermediate metallicities.

Chapter 5. Systematic Effects on Oxygen Abundances Derived from the 6300 Å [O I] and 7774 Å O I Lines

Abstract

We consider potential systematic effects on Oxygen abundances of F and G dwarfs derived from the 6300 Å [O I] line and the 7774 Å O I triplet. Our observed Solar intensity spectra of the 7774 Å triplet confirm Altrock's (1968) finding of a discrepancy of observed equivalent widths compared to both LTE and recent NLTE models. However, we emphasize that the disagreement at low μ does not seem to affect the Solar O abundance as derived from flux spectra since such values are in excellent agreement with the meteoritic value. The Solar 7774 Å equivalent widths from flux spectra are equal to those from intensity spectra and LTE models at $\mu = 0.61-0.66$, the value expected from the classic Eddington-Barbier relation. Since we know of no reason this relation should break down for more metal-poor stars, it seems reasonable that LTE 7774 Å abundances for metal-poor dwarfs should be just as accurate as in the Solar case, all other things being equal (*i.e.*, assuming that NLTE effects and thermal inhomogeneities are not significant in more metal-poor atmospheres). We note that the 6300 Å equivalent width for the Sun seems to be rather uncertain. Because of this uncertainty and the inability of authors to reproduce each others' 6300 Å O abundances, we do not believe it can be claimed *a priori* that these abundances are more reliable than those derived from the 7774 Å triplet. This would seem especially true in light of recent pseudo-two-stream calculations, which indicate that the effects of atmospheric inhomogeneities on derived O abundances may be *larger* for the 6300 Å [O I] line. Hence, O abundances from the forbidden line may be subject to systematic uncertainties of the same magnitude as claimed for the permitted triplet. We empirically investigate the situation by deriving abundances for a selection of relatively metal-rich F and G dwarfs. For $T_{\text{eff}} \gtrsim 6200-6300$ K, 7774 Å abundances are substantially larger than the 6300 Å abundances. This probably reflects the onset of well-established NLTE effects for early-type stars. For $T_{\text{eff}} \lesssim 6200-6300$ K, however, we find no systematic difference between the 6300 and 7774 Å abundances. This result is in conflict with that of Nissen & Edvardsson (1992), whose data we reanalyze here. The discrepancy likely resides in the model atmospheres used in the analysis. *If* systematic differences are present, then our analysis seems to have successfully removed them. If recently proposed, hotter T_{eff} values for metal-poor dwarfs are correct, then there appears to be no discrepancy between the 6300 Å abundances of metal-poor giants and the 7774 Å abundances for dwarfs. This would seem to rule out the presence of substantial NLTE departures (in agreement with the computations of Tomkin

et al. 1992). Small effects (*i.e.*, $\lesssim 0.05$ dex say) due to LTE departures and thermal inhomogeneities, however, cannot be excluded. While more observational and theoretical work remains, we conclude that 7774 Å O abundances may be just as reliable as those from the 6300 Å line in both metal-poor and metal-rich dwarfs (for $T_{\text{eff}} \lesssim 6200 - 6300$ K of course).

1. Introduction

The O abundance (and O/Fe ratio) of stellar atmospheres/interiors is an important parameter. O abundances provide both important clues concerning Galactic chemical evolution (Wheeler, Sneden & Truran 1989) and constraints on galaxy formation scenarios (Gilmore, Wyse & Kuijken 1989, Wyse & Gilmore 1992). Additionally, stellar evolutionary models can be affected by the choice of the [O/Fe] ratio. Such models are used to produce isochrones from which important quantities such as globular cluster ages are derived. Unfortunately, stellar O abundances (particularly in the halo) are, currently, rather uncertain. Analysis of the 6300 Å [O I] line in halo *giants* indicates that [O/Fe] $\sim +0.4$ to $+0.5$ (*e.g.*, Barbuy & Erdelyi-Mendes 1989; Barbuy 1990; the reanalysis of Bessel, Sutherland & Ruan 1991). Abia & Rebolo (1989) determined O abundances in halo *dwarfs* using the 7774 Å O I triplet. Their surprising results indicated that [O/Fe] increased with decreasing metallicity, reaching $\sim +1.2$ for stars with $[\text{Fe}/\text{H}] \leq -2.5$.

It has been repeatedly suggested (Kiselman 1991; Spiesman & Wallerstein 1991; Spite & Spite 1991; Nissen & Edvardsson 1992, hereafter NE92) that the large discrepancy in the [O/Fe] ratios of metal-poor giants and dwarfs could be caused by either (or both): a) NLTE effects in the analysis of the high-excitation 7774 Å O I triplet b) Thermal inhomogeneities, again affecting the 7774 Å abundances, which are not properly accounted for in the model atmospheres. Some evidence that systematic effects in the 7774 Å lines are at play comes from the analysis of the 6300 Å [O I] line (which is not thought to be subject to the effects of NLTE departures or thermal inhomogeneities; Altrock 1968, Lambert 1978, Kiselman 1991) in halo dwarfs. Spiesman & Wallerstein (1991) found [O/Fe] = $+0.3$ and $+0.4$ for two dwarfs having $[\text{Fe}/\text{H}] = -1.74$ and -1.43 . Spite & Spite (1991) found [O/Fe] = $+0.48$ and $+0.59$ for two dwarfs having $[\text{Fe}/\text{H}] = -1.4$ and -1.6 . These [O/Fe] values are significantly lower than the values derived by Abia & Rebolo (1989) for halo dwarfs. NE92, in their analysis of 23 F and G dwarfs, also found that the O abundances derived from the permitted triplet are larger than those derived from the forbidden line (by ~ 0.2 dex at $[\text{Fe}/\text{H}] = -0.8$) and that this difference increases with decreasing metallicity.

Kiselman (1991) theoretically investigated NLTE effects on O abundances derived from the

7774 Å triplet. Comparison of the top and middle panels of his figure 4 indicates that his NLTE calculations are able to reduce the [O/Fe] ratios of the most metal-poor stars by ~ 0.4 dex on average, bringing them in closer agreement with the ratios derived using the 6300 Å [O I] line. However, the solar disk equivalent widths of the 7774 Å O I lines predicted by Kiselman's (1991) NLTE models are in gross conflict with the observations of Altrock (1968). Kiselman's (1991) LTE calculations are in good agreement with the solar disk center equivalent widths; however, the variation of the equivalent widths with the usual cosine angle, μ , across the solar disk is better reproduced by the NLTE models than by the LTE models. This confusing situation is exacerbated by the NLTE calculations of Abia & Rebolo (1989) and Tomkin *et al.* (1992). These authors' calculations indicate that NLTE effects on O abundances derived from the 7774 Å O I triplet are very small (amounting to a few hundredths of a dex). The difference between Kiselman's results and those of Tomkin *et al.* seems to be due to the latter's inclusion of collisions with neutral H atoms, which drives line formation back into LTE.

In chapter 1 (King 1993a), an explanation was proposed for the larger O abundances derived from the 7774 Å permitted triplet in metal-poor stars. This explanation contained two basic parts: (1) The permitted O I triplet equivalent widths of Abia & Rebolo (1989) appear to be systematically too high (a result also found by Tomkin *et al.* 1992) and (2) Current temperature estimates for metal-poor stars may be systematically too low. After accounting for these two factors, the mean [O/Fe] level derived from observations of the 7774 Å triplet in metal-poor dwarfs ([O/Fe] = +0.53) was found to be in good agreement with the value deduced from the forbidden line in halo giants ([O/Fe] = +0.46, see chapter 4 or King 1993b). The remaining small (~ 0.07 dex) difference between these mean [O/Fe] ratios does not appear to be significant in light of the observational uncertainties. However, there is, of course, a possibility that the small difference could be due to NLTE departures (or thermal inhomogeneities) in the 7774 Å lines. The size and magnitude are not totally inconsistent with the calculations of Tomkin *et al.* (1992).

The purpose of this chapter is to address some questions concerning O abundances derived from the 7774 Å permitted triplet *versus* those derived from the 6300 Å forbidden line: (1) For the Sun, the LTE models do not predict the correct center-to-limb variation of the 7774 Å triplet equivalent widths. Are observational data of Altrock (1968), used in the comparison of Kiselman (1991), correct? If so, what are the implications for determining O abundances from the 7774 Å triplet? (2) Can the small effects of NLTE departures predicted by Tomkin *et al.* (1992) be reconciled with the larger differences between the 6300 Å and 7774 Å abundances inferred by NE92? Can the small difference between the 6300 Å and 7774 Å abundances that remains after adopting revised T_{eff} values from chapter 2 be reconciled with the larger differences inferred by NE92? (3) The effects of

NLTE departures on O abundances derived from the 7774 Å lines in early type stars are very large and seem unquestionable (*e.g.*, Baschek, Scholz & Sedlmayer 1977). It is thus expected that the difference between the 6300 Å and 7774 Å abundances will be an increasing function of increasing T_{eff} as NLTE calculations indeed predict (Kiselman 1991). The next chapter will discuss O abundances in open cluster stars. Most of the cluster dwarfs or subgiants are much fainter than the field disk stars having had O lines observed to date. This circumstance could potentially introduce a bias in the O abundance analysis via two avenues. First, one might observe brighter (hence hotter) main-sequence cluster stars to achieve larger S/N data. Second, in observations obtained with a multi-object spectrograph, the highest S/N spectra will be the brighter (hence hotter) objects. If one restricts the analysis to the hotter main-sequence stars in open clusters, severe NLTE effects may enter the analysis. It must then be determined, in advance, at what temperatures these NLTE effects become significant. Then, hotter stars can be excluded from (or possibly have their abundances corrected during) the analysis. We address these three issues here.

2. The Solar Center-to-Limb Variation of the 7774 Å O I triplet

2.1 Observations and Data Reduction

Solar observations of the 7774 Å O I triplet were conducted on February 17, 1993 at Mees Solar Observatory on Haleakala, Maui. The data were obtained with the Mees CCD (MCCD) imaging spectrograph which has been described by Penn *et al.* (1991). Briefly, the instrument is a spatially-scanning imaging spectrograph consisting of the 0.25-m coronagraph telescope, a 3.0-m coude spectrograph, and a rapid-writing data storage system. A 300 l/mm echelle grating was used in 7th order with an order sorting filter having a central wavelength of 7757 Å and a band width of 115 Å. This configuration gave a dispersion of 0.021 Å per pixel. The resolution was set by adjusting the slit opening. Unfortunately, while adjusting the slit, the device used to measure the width ceased to function. Our best estimate was a slit width of ~ 2.4 pixels and we know for a fact it was ≤ 2.6 pixels. The detector was a cooled, UV-coated 384×576 Thompson CCD with 22 μm pixels.

The first step in the observations was to locate Sun center. This was done by monitoring the Solar image from a video rate CCD camera and associated instrumental encoder readings. Centering was performed in both the N-S and E-W directions. The position of the solar limb in our series of images was in excellent agreement with that expected from our centering procedure. This assures us that the sequence of scans across the disk was made along a projection of a great circle rather than along a chord whose origin was displaced from disk center.

Two 1-second scans were taken at each position, beginning at disk center, then moving south to the limb (with the instrument scanning in the E-W direction at each position), and back to disk center. Four pixel binning was employed in one of the spatial directions. Calibration frames (biases, darks, and flats) were taken before and after the entire sequence of scans. Five 1-second scans at disk center were also obtained for a wavelength region in the midst of the atmospheric (O_2) A-band near 7600 Å. This was done as a check on scattered light in the spectrograph. The same instrumentation was employed except for the use of an order sorting filter centered at 7610 Å with a band width of 360 Å.

The reductions began with removal of the dc-offset via a mean bias frame and the removal of dark current via a mean dark frame. The next step in the reductions is usually the flat-fielding of the object frames to remove pixel-to-pixel variations. Unfortunately, these frames (taken by an automated calibration sequence optimized for a different, long-term program at Mees) were not sufficiently exposed to be useful. The Thompson CCD used is known to be flat at a level of somewhat less than 1%. This was confirmed by our measured final S/N values which are typically ~ 130 in the continuum near the O I triplet. Even with this relatively low S/N, the predicted 1σ uncertainty (Cayrel 1988) in the measured equivalent widths is ± 0.6 mÅ given our resolution. Measurements of the equivalent widths from spectra taken at the same position but in the two separate scans indicate the real value is slightly larger than ~ 1.0 mÅ. This is certainly more than adequate for a comparison with the equivalent widths of Altrock (1968). Three spatial (N-S) elements were collapsed and summed in the data cube to produce a final spectrum for each scan. The spectra measured were always taken from the same area of the data cube for the sake of consistency. No summing in the other spatial dimension was performed as the spectral lines had a significant tilt in the spatial *vs.* spectral plane. Wavelength calibration was performed using the wavelengths of the O I triplet lines themselves as well as that of the strong Fe I line at 7780 Å. The rms residuals of the wavelength solutions were ~ 0.003 Å.

2.2 Results

Two representative spectra at different disk positions are shown in Figure 22. Equivalent widths of the O I lines were measured in the IRAF package. Values from the two different scans at the same disk position were averaged to yield the values listed in Table 9. Our results are contained in Figures 23a-23c which compare our results (filled symbols) with those of Altrock (1968; open symbols) for each line in the triplet. The solid lines are LTE model values calculated for an O abundance of log

$\epsilon(\text{O}) = 8.91$ (consistent with the meteoritic value of 8.93, Grevesse & Anders 1989) using the new Kurucz (1992) solar model. Figure 23d is analogous to figure 2 of Kiselman (1991), showing the data of Altrock (points) for each line in the triplet and the LTE model calculations (solid lines). Figure 24 shows the data of Altrock (1968) compared to a solar NLTE model kindly provided by Drs. J. Tomkin and M. Lemke. A description of their model can be found in Tomkin *et al.* (1992).

2.3 Discussion

2.3.1 Comparison of our data with Altrock

Figures 23a–23c indicate that significant discrepancies exist between our data and that of Altrock (1968), particularly at larger values of μ . Both have typical uncertainties of $\sim 1 \text{ m\AA}$ for each datum, so the disagreement is, at first, difficult to understand. The most plausible explanation is the presence of scattered light, having an increasing effect as one proceeds to the brighter disk center, in our data. Scattered light would decrease our equivalent widths relative to Altrock’s as is observed.

Some effort was made to measure the scattered light in the spectrograph by obtaining observations of the very strong atmospheric A band near $7600 \text{ \AA}$ at disk center. In the Kurucz *et al.* (1984, hereafter KFBT) atlas, the cores of these telluric lines reach the zero flux level. Adopting a very localized continuum makes the conservative (but poor) assumption that the peaks between the lines represent true continuum. For this case, the A-band cores are $\sim 6.7\%$ above zero flux. A continuum level set by the region just to the blue of the A-band results in a value of 4.7% . Using the KFBT atlas for guidance in setting the continuum also yields a scattered light contribution of 4.7% . This level of scattered light can account for the discrepancy between our disk-center measurements and those of Altrock (1968).

Solar “surface” features (*e.g.*, spots) were readily evident in both the spatial-spatial and spatial-spectral plane. We are confident that our final spectra avoid such regions. As mentioned before, we are also confident that our sequence of scans was not taken along a chord extending from off disk center to the solar limb. Although we can offer no explanation for any remaining differences between our measurements and Altrock’s, we hold his data in high regard because of his use of a double-pass spectrograph, which minimized scattered light contributions to 1% . In any case, despite these differences, our data still serve to confirm the difference between the measured Solar O I equivalent widths and the results of LTE calculations at small μ .

2.3.2 What ARE the Solar O Equivalent Widths?

The disk center equivalent widths of the O I triplet lines might be regarded as somewhat uncertain, judging by comparison of independent measurements. Table 10 presents various estimates. The mean value and variance of the disk center measurements of Altrock (1968), the Liege atlas, and the Utrecht atlas, seems to indicate that our *scattered light corrected* values cannot be considered discrepant. Integrated (flux value) equivalent width measurements are presented in the bottom half of Table 2. The value from the KFBT atlas seems a bit larger than more recent measurements of E93, AMB (using the asteroid Vesta), and JRK (bright sky spectrum). The values used by NE92, E93, and those measured by AMB and JRK from high and low dispersion spectra are in good agreement.

We note that the situation for the 6300 Å line may be worse. Lambert (1978) adopted 4.55 mÅ for the solar equivalent width (N. B. this is a disk center value). NE92 adopt a value of 5.3 mÅ based on the KFBT atlas though they measure a value of 5.7 mÅ from their own spectrograms. The 6300 Å equivalent width seems to be very sensitive to the adopted continuum. We also measured equivalent widths using the KFBT atlas. Adopting a pseudo-continuum just above apparent continuum “windows” yielded 5.4 mÅ. Using a continuum setting that we applied to our stars resulted in 5.9 mÅ. Using the continuum normalization of KFBT, however, results in ~ 7.9 mÅ. For consistency, we will later use the same continuum level applied to our stellar spectra. However, it seems clear that the solar equivalent width can potentially be a source of significant systematic differences in stellar [O/H] ratios.

2.3.3 The slope of the equivalent width— μ relation

An important conclusion we draw from our data, despite the effects of scattered light near disk-center, is the marked flatness of the equivalent width — μ relation as compared to the LTE models. This departure between our data and the LTE model can be seen in figures 23a–23c and in figure 23d for Altrock’s (1968) data. Kiselman’s (1991) NLTE calculations match well the observed shape of the variation with μ ; however, his figure 2 shows that the magnitude of his disk-center equivalent widths are much too large. Tomkin *et al.* (1992) improved upon Kiselman’s (1991) calculations by including excitation effects due to collisions with neutral hydrogen. Figure 24 shows their NLTE calculations (solid lines) *versus* Altrock’s (1968) data. Clearly, the agreement is much better. Indeed, Tomkin *et al.* (1992) vary the collision rates in order to force better agreement.

It is interesting to compare the equivalent widths of various flux and intensity measurements. Comparing the mean of the flux measurements by Edvardsson *et al.* (1992) and AMB with the

intensity measurements of Altrock (1968) and with the LTE model calculations in Figure 24, one finds that the flux values are equal to the intensity measurements at $\mu = 0.61 - 0.66$. This is quite close to what one expects by adopting the classical Eddington-Barbier relation (where the source function is a linear function of optical depth). Figures 23a–23c show that the LTE models are in good agreement with Altrock’s data at these values of μ . If the same situation holds for halo dwarfs, then we can argue that LTE analysis should yield correct abundances for these stars too (assuming NLTE effects and thermal inhomogeneities are not prevalent, since the fact that the data conform to the Eddington-Barbier relation say nothing about the validity of the LTE assumption)¹⁰. We probably expect Eddington-Barbier to be a good approximation in metal-poor stars where the O I triplet lines are weaker than the Sun’s, lying securely on the linear portion of the curve of growth, and do not exhibit substantial wings. In contrary cases, the source function is a higher order function of the optical depth and Eddington-Barbier assumptions may not be valid (*i.e.*, in opposite cases, the line profile outside the Doppler core may be a rapidly varying function of frequency or, for a strong line, there may be a large gradient [with height] in the Doppler width). In any case, we re-emphasize that the Eddington-Barbier relation says nothing about LTE formation. The point to be stressed is that LTE analyses of Solar 7774 Å O I triplet flux measurements seem to yield the “correct” (as taken from the meteoritic value) Solar abundance. Whether or not LTE analyses yield “correct” results for other stars, depends on NLTE departures and the extent of atmospheric inhomogeneities in these stars.

3. 6300 and 7774 Å O Abundances in F and G Dwarfs

¹⁰ It must be emphasized at this point that we do *not* deny that the 7774 Å lines likely form in NLTE conditions for these stars. Indeed, it is well-known (*e.g.*, Altrock 1968) that the triplet forms in NLTE even for the Sun. What we are concerned with here is if the NLTE formation has a significant effect on abundances derived from *flux* spectra. For the Sun, this is certainly not the case as we will later find a solar abundance of $\log \epsilon(\text{O}) = 8.89$ from our flux spectra. This value is perfectly consistent with the meteoritic abundance of 8.93 (note that our abundance derived from flux spectra is *less* than meteoritic and thus does *not* reflect a real systematic difference due to NLTE formation, which would lead to a *higher* abundance). In this chapter, “NLTE effects” will refer to departures which result in significant effects on stellar abundances derived from flux spectra. Nor do we deny that inhomogeneities *are* present in the solar atmosphere. Indeed, the presence of inhomogeneities can be inferred from the asymmetries, seen in the O I triplet, which vary as a function of μ . These asymmetries were noted by Altrock (1968), Lambert (1978), and can be seen in our data.

We now compare the O abundances of F and G dwarfs derived from the 6300 Å [O I] line and the 7774 Å O I triplet. Such a comparison allows us to determine empirically if systematic effects are present in the analyses and if such effects are functions of [Fe/H] and/or T_{eff} . We combine our own data taken for this purpose with data from the literature which we reanalyze for homogeneity,

3.1 The Data

Observations of the 6300 Å [O I] line and 7774 Å O I triplet in F field dwarfs were obtained by Dr. A. M. Boesgaard using the 3.6-m Canada-France-Hawaii telescope during observing runs in August and October 1987. The instrumentation consisted of the coudé spectrograph (including the red mirror train), a 830 l/mm grating, the f/8.2 camera, and a cooled 1872 Reticon detector (15 μm pixels). This configuration yielded a dispersion of 0.067 Å/pix while the measured resolution of the spectra was 0.11 Å. Rapidly rotating B stars were observed at airmasses similar to the program objects in order to facilitate the removal of telluric lines which are present in the 6300 Å region. Comparison spectra were obtained each night in order to wavelength calibrate the reduced spectra. Multiple series of four flat-field spectra were obtained each night. Each set of four flat fields was iteratively summed and averaged with the previous set of flat fields. Then, each stellar spectrum was divided by the averaged flat field whose mean was within 10% of the stellar exposure level. In addition, a 4-channel normalization was done to remove remaining pixel to pixel variations of the detector. These flattened spectra were then placed on a wavelength scale, continuum normalized, and divided by the spectra of the rapidly rotating B stars (which were reduced in the same way) in order to remove telluric features. The S/N of the final spectra are typically ≥ 400 .

Observations of the 6300 Å and 7774 Å O lines in additional F and G dwarfs were also obtained by JRK on the nights of August 23–24 and 25–26, 1991 using the coudé spectrograph of the University of Hawaii 2.2-m telescope on Mauna Kea. The detector used was a 1024×1024 Tektronix CCD. The measured resolution of the spectra was 0.16 Å. A blocking filter was used to prevent contamination from higher orders. Exposure times ranged from 300 to 700 seconds for the 6300 Å observations and from 700 to 1200 seconds for the 7774 Å observations. The S/N of the reduced spectra is typically 300–400.

Standard reductions were carried out on the raw data. This included trimming the CCD frames and removing the dc-offset of the detector by subtracting a mean value from the overscan region. The flat field system at the 2.2-m coudé is rather rudimentary, consisting of an incandescent bulb illuminating an ordinary sheet of 8.5"×11" paper which, in turn, illuminates the slit. Approximately

7 flat field frames were bias-subtracted and averaged (using 3σ clipping) to form a nightly flat field frame. This frame was then fit with a low order cubic spline in the direction of dispersion to remove large-scale structure. Tests show that we are able to flatten our data to the $\lesssim 0.4\%$ (per pixel) level. The bias-subtracted data were divided by the nightly flats in order to remove both large-scale structure and the small-scale pixel-to-pixel variations. A low order cubic spline was fit in the dispersion direction some 10-20 pixels adjacent to (but well outside the wings) of the spatial profile of the stellar spectrum in order to remove residual background (the small dark current, sky background, bias levels slightly elevated above that in the overscan region, and extraneous light). The spectra were then traced and extracted using IRAF. Wavelength calibration was performed using 5-10 lines from Th-Ar lamp spectra. The wavelength solutions utilized a low order Legendre polynomial and the final solutions have rms residuals of ~ 0.005 Å. The wavelength-calibrated spectra were then continuum normalized. Finally, the 6300 Å spectra were divided by spectra of rapidly-rotating, early-type stars to remove telluric lines. A typical spectrum of the 7774 Å region is shown in Figure 25.

Both 6300 and 7774 Å O data were also taken from the literature to supplement our own. NE92 have derived O abundances for 23 F and G dwarfs from these lines. We rederive abundances in a manner consistent with our own data. We also reanalyze the data of Clegg, Lambert & Tomkin (1981). All stars included in our analysis are listed in the first column of Tables 11 and 12.

3.2 Stellar Parameters

3.2.1 [Fe/H]

Although iron abundances were derived by us for a few of our stars, we supplement these estimates with recent values from the literature. Edvardsson *et al.* (1993, hereafter E93) have derived [Fe/H] for a large fraction of our stars. We also utilize the [Fe/H] values derived by Balachandran (1990). These values are combined with unpublished determinations by AMB and with other data from the literature (we relied heavily on the catalog of Cayrel de Strobel *et al.* 1992) to determine mean [Fe/H] values which are listed in column 6 of Table 3. The agreement between various sources is very good. We estimate a typical uncertainty of $\pm 0.10 - 0.15$ dex in our adopted values.

3.2.2 T_{eff}

For most of the stars, T_{eff} values were determined from the Saxner & Hammarback (1985) relations. $(B - V)$ colors were taken from the Bright Star Catalog (Hoffleit & Jaschek 1982) or

the SIMBAD database. Stromgren indices, $(b - y)$ and $H\beta$, were taken from Hauck & Mermilliod (1990). The photometric indices and the adopted $[\text{Fe}/\text{H}]$ values were used to determine mean T_{eff} estimates using the Saxner & Hammarback (1985) relations. The final values are listed in column two of Table 11. Uncertainties in the T_{eff} values, based on the variance of the determinations from each photometric index, are listed in column three. For the five most metal-poor stars studied by NE92, T_{eff} values were derived using the $(b - y) - T_{\text{eff}}$ relation given in chapter 2. These T_{eff} values are only slightly larger than those adopted by NE92. The photometric indices of HR 7462 lie outside the range of the Saxner & Hammarback (1985) calibrations. We derived the T_{eff} for this star by mating the theoretical $(B - V)$ - and $(b - y) - T_{\text{eff}}$ relations of the new Kurucz (1992) models to the Saxner & Hammarback (1985) relations at $(B - V) = 0.63$.

Our temperatures are, on average, slightly lower than those given by E93. For 38 stars we have in common, the mean difference (in the sense of our temperatures minus those of E93) is -36.2 ± 38.5 K. Plotting the differences versus our T_{eff} values reveals a trend in the sense that we measure smaller T_{eff} values than E93 for hotter stars. The trend is small and appreciable scatter is present, but the ordinary correlation coefficient indicates that the effect is significant at the 99.5% confidence level. We note that a more rigorous parameter determination would involve re-solving for the metal abundances using the T_{eff} adopted here. This modified $[\text{Fe}/\text{H}]$ value would then be used to recalculate T_{eff} using the Saxner & Hammarback (1985) relations and the process iterated until convergence is attained. However, in our case, the changes are extremely small and well within the estimated uncertainties in the parameters. Therefore, we retain the $[\text{Fe}/\text{H}]$ and T_{eff} values initially estimated.

3.2.3 $\log g$

Two major means were used to estimate the gravities of the stars in our sample. First, when available, we took the estimates from E93. These are based on a Stromgren photometric calibration. Second, we computed the gravities based on the $(b - y)$ and c_1 indices of the new Kurucz (1992) models. Curves of constant gravity were plotted in the c_1 *versus* $(b - y)$ plane for models of different metallicity. The value of $\log g$ was then read off these curves. The difference between the two determinations (in the sense E93 minus Kurucz) is plotted against the Kurucz model $\log g$ value in Figure 26. A marked trend in the differences is evident. The Kurucz model values were placed on the scale of E93 (which we feel has greater external accuracy) via the relation for the differences

$$\Delta \log g = -0.662 \times \log g(\text{Kurucz}) + 2.908$$

This relation is shown by the solid line in Figure 26. The transformed Kurucz values were averaged with the E93 values to increase the internal accuracy of the determinations. We also drew on gravity estimations from the literature for a few of the stars. These were used, in all but a few cases, only to estimate uncertainties. Our final $\log g$ values are shown in column four of Table 3. The estimated uncertainties are shown in column five.

3.2.4 Microturbulent Velocity

Microturbulent velocities were derived via equation (9) in E93. This relation was derived empirically and is similar to that found by Nissen (1981). Our values, listed in the final column of Table 3, are nearly identical to those used by E93.

3.3 Equivalent Widths

The IRAF and DOUBLEPLAY packages were used to measure the equivalent widths from our spectra. Several direct integration and Gaussian fits were used for both the 6300 Å and the 7774 Å lines to provide values and uncertainties. Table 12 lists the measured 6300 Å equivalent width and its uncertainty in columns 2 and 3. The uncertainties for NE92’s equivalent widths represent the minimum-maximum variation of their individual measurements. For other literature values, we assigned reasonable (but necessarily subjective) uncertainties based on resolution and S/N. Sources of the 6300 Å values are given in column 4 of Table 12. Columns 5 and 6 give the equivalent width and uncertainty of the *whole* 7774 Å triplet. Column 7 lists the source of the data. Again, we were forced to assign uncertainties for some values. NE92 do not provide 7774 Å equivalent widths so no values are given. In the next section we describe how 7774 Å O abundances for these stars were determined.

Table 12 also gives equivalent width data from multiple sources. However, only one source was used to derive O abundances. We preferred the 6300 Å measurements of AMB and CLT over those of JRK, due to the latter’s lower resolution. For the sake of homogeneity, the same source was used for both the 6300 Å and 7774 Å measurements. We do make abundance estimates from both listed sources for HR 4540 because doing so allows a comparison between our direct abundance determination made from the CLT equivalent widths and our indirect reanalysis of the NE92 data. The two determinations are treated as independent data points in the ensuing discussion. Comparison of multiple measurements indicates very good agreement between the 6300 Å and 7774 Å equivalent widths of AMB, JRK, and CLT.

We initially treat the Sun as an individual data point in the analysis. The solar 7774 Å equivalent width was measured from a high resolution (0.21 Å) spectrum, acquired by AMB, of the asteroid Vesta. We set the continuum in a manner consistent with the stellar spectra. The solar 6300 Å [O I] line was measured from the KFBT atlas. An average of the equivalent widths derived from both direct integration and Gaussian fits was adopted. We did not use the continuum level of KFBT, but instead used the same continuum setting as for our stellar spectra. The equivalent widths and estimated uncertainties are given in Table 12. We will normalize our stellar abundances to Solar later. This normalization makes the stellar abundances independent of the adopted gf values and, to first order, the adopted continuum fit.

3.4 Abundance determinations

O abundances were determined from the measured equivalent widths using a curve of growth analysis. We use the LTE analysis software, RAI10, provided by Dr. M. Spite and the new model atmospheres of Kurucz (1992) to construct the curves of growth. Models from Kurucz’s (1992) extensive grid were chosen according to the T_{eff} , $\log g$, and [Fe/H] values of the individual stars. Due to the large number of curves of growth constructed, it was an easy matter to interpolate among them (using low order polynomials) to determine the abundance for a certain equivalent width and set of stellar parameters.

3.4.1 6300 Å [O I] Abundances

We adopted $\log gf = -9.75$ for the 6300 Å [O I] line (Lambert 1978). The same value was used by NE92. The curves of growth computed for each star yield an abundance given the measured equivalent width. Column eight of Table 12 gives the abundance, $\log \epsilon(\text{O})$, from the 6300 Å line. Because we treat the Sun as an individual point for the moment, the abundances are expressed on the usual logarithmic scale with $\log \epsilon(\text{H}) = 12.0$. Uncertainties in the 6300 Å abundance, determined by adding the uncertainties due to errors in the equivalent width; T_{eff} ; $\log g$; and [Fe/H] in quadrature, are listed in column nine of Table 12.

We compare the absolute O abundances, $\log \epsilon(\text{O})$ and *not* [O/H], of NE92 with those found in our reanalysis. The mean difference (K93–NE92) is -0.010 ± 0.047 (1σ) dex. The small difference is precisely that expected from the slightly lower, on average, T_{eff} values adopted here. The excellent agreement we find indicates that the different software and model atmospheres utilized in the analysis do not significantly affect our results– at least for the 6300 Å abundances.

3.4.2 7774 Å O I Abundances

We adopt $\log gf = +0.333$, $+0.186$, and -0.035 for the 7772, 7774, and 7775 Å O I lines (Wiese, Smith & Glennon 1966). These values are significantly different from those of NE92. Our gf values are ~ 0.105 dex larger, on average, than those used by NE92 (kindly communicated to us by Dr. P. Nissen). This difference represents the first main discrepancy we find between our results and those of NE92—namely, we are unable to reproduce their astrophysically derived 7774 Å gf values.

Two discrepancies appear by illustrating the situation in the following steps: 1) NE92 adopt a value for the solar O abundance of 8.90. We determined a solar O abundance using our adopted gf values, but we can also work backwards and describe what gf values would be derived using our O abundance. We “adopt” our derived solar O abundance of 8.89. 2) We use a 7774 Å triplet equivalent width of 179.2 mÅ. NE92 use a value of 175.8 mÅ (Dr. P. Nissen 1993, private communication). E93 determine O abundances which are predicated, in part, on the results of NE92. E93’s gf values, as determined from an assumed solar abundance of 8.93 and an equivalent width of 176.8 mÅ, are 0.03 dex *larger* than those of NE92 despite the fact that a *larger* solar O abundance (8.93 *versus* 8.90) was assumed by the former. This is opposite to what one expects, all other things being equal, since the quantity $\log Ngf$ is constant in an LTE analysis. The difference in $\log Ngf$ could be explained if NE92’s equivalent widths were somewhat larger than those of E93. However, as noted, the values are actually lower (though insignificantly so). This is the first discrepancy. 3) Both we and NE92 use essentially the same solar equivalent width (176 *versus* 179 mÅ). Both we and NE92 use essentially the same solar O abundance (8.89 *versus* 8.90). Yet, we would find gf values differing by 0.10 – 0.11 dex. This is the second discrepancy. We can offer no definitive explanation for either discrepancy. Regarding the first discrepancy, however, Dr. Nissen has informed us that their continuous opacity code was updated between the analysis of NE92 and E93. We do not know if this can account for the noted difference.

We used a radiative damping constant of 1.3×10^8 . The classical van der Waals damping was calculated according to the formulation of Unsold (1955), using the expression for C_6 from Gray (1992). An enhancement factor of ~ 1.2 was adopted. E93 use slightly different values for these parameters and so we assume that this was also done in NE92. We constructed curves of growth for a range of T_{eff} , $\log g$, and $[\text{Fe}/\text{H}]$ using the E93 constants. Differences in the abundances (or $\log gf$ values) always amounted to ≤ 0.01 dex (and usually about half this value). Hence, the discrepancy in the gf values is unresolved. We did note that the different software and model atmospheres used by

us and NE92 resulted in very good agreement for the 6300 Å abundances. Because the 7774 Å lines are formed deeper in the stellar photosphere, a likely cause of the discrepancy may be differences in the model atmospheres used in the analyses. NE92 employed new, updated Gustafsson (Gustafsson *et al.* 1975) atmospheres.

Since 7774 Å equivalent width measurements were not reported by NE92, we rederive abundances as follows. First, we computed the absolute 7774 Å abundances found by NE92 using the values in columns six and two of their Table 4 and by adopting their assumed solar abundance of 8.90. Then we constructed curves of growth, using our own adopted stellar parameters. Adjustments to their absolute abundance were made by examining the sensitivity of the abundances to T_{eff} , $\log g$, ξ , etc. from the curves of growth. Finally, the revised absolute abundances were reduced by 0.105 dex to place them on our gf scale. Our 7774 Å abundances and uncertainties for both the NE92 stars and the remaining stars (for which abundances were derived directly using their equivalent widths), are given in columns 10 and 11 of Table 12. Because HR 8181 had 6300 Å data from a source other than four main ones we used, it was of interest to locate a 7774 Å measurement. The only data are those of E93. Thus, this star’s abundance was rederived using the same procedure as for the NE92 data. First, though, we converted the “corrected” O abundance of E93 to a 7774 Å abundance using their equation (11).

We now compare the *absolute* 7774 Å abundances of NE92 to those of this reanalysis. The mean difference (K93–NE92) is -0.093 ± 0.033 dex. As noted above, our reanalyzed 7774 Å abundances of NE92 have been corrected for our choice of $\log gf$. If this were not done, the mean difference would be $\sim +0.012$ dex. This value is exactly what is expected since our T_{eff} values are, on average, slightly lower than those of NE92. We also note at this point that abundances were derived from the total equivalent width of the triplet, rather than averaging the abundances derived from the individual lines. It was found that the latter method did not yield different abundances. Indeed, Boesgaard & King (1993) found that their O abundances derived from the individual lines of the triplet are in very good agreement and that the average of these values is not significantly different than the value derived from the triplet as a whole. This result was confirmed here since one of us (AMB) had previously derived abundances from the individual lines and found them to be in excellent agreement. Additional tests show that there is no significant difference between the abundances derived from the average of the three triplet lines and the triplet as a whole.

3.5 Results and Discussion

3.5.1 $[O/H](6300 \text{ \AA})$ vs. $[O/H](7774 \text{ \AA})$

We take our absolute O abundances and normalize them to the appropriate (6300 Å or 7774 Å) solar abundance we derived (given in Table 12). The O abundance (relative to Solar) from the 6300 Å [O I] line is plotted against the O abundance from the 7774 Å triplet in Figure 27. The open squares are data points from the NE92 sample while the filled squares represent all other data. The dashed line represents the relation expected if the abundances from the permitted and forbidden lines are equal. It can be seen that, in general, the [O/H] ratios as determined from the 7774 Å triplet are larger than those derived from the 6300 Å line. Below, we will find that this behavior is due to the effects of temperature and the presence of systematic differences between the NE92 data and the other data.

3.5.2 $\Delta[O/H]$ vs. T_{eff}

Taking our appropriately normalized O abundances, we calculate the quantity $\Delta[O/H] = [O/H](7774) - [O/H](6300)$. $\Delta[O/H]$ is plotted *versus* T_{eff} in Figures 28a–28c for the reanalyzed data of NE92, all other data, and both samples respectively. Figure 28c clearly shows a large increase in $\Delta[O/H]$ for $T_{\text{eff}} \gtrsim 6200 - 6300$ K. This is expected since NLTE effects on 7774 Å abundances are known to exist in early-type stars (Baschek *et al.* 1977). It seems certain that we are seeing the onset of these NLTE effects. Thus, when we analyze O abundances of open clusters in chapter 6, we restrict our attention to stars cooler than $\sim 6200 - 6300$ K.

Comparison of Figure 28a and 28b reveals the second main discrepancy, noted earlier, between this work and that of NE92. For the reanalyzed NE92 stars having $T_{\text{eff}} \leq 6200$ K, the mean value of $\Delta[O/H]$ is $+0.092 \pm 0.086$ (1σ) dex. The mean for the rest of our sample having $T_{\text{eff}} \leq 6200$ K is 0.026 ± 0.094 dex. The higher $\Delta[O/H]$ ratios we find for the NE92 stars are independent of the adopted gf values and solar abundance. This is demonstrated in Figure 29, which shows the original data of NE92; these data have not been adjusted in any way. The mean $\Delta[O/H]$ for stars having $T_{\text{eff}} \leq 6200$ K in this case is $+0.117 \pm 0.094$.

NE92 took their results as evidence that systematic differences existed between O abundances derived from the 6300 Å and 7774 Å lines for their sample of relatively metal-rich stars. Restricting ourselves to stars with $T_{\text{eff}} \leq 6200$ K, we find that our data, shown in Figure 28b, are not consistent with this conclusion. Rather, we are not able to find any significant difference between the [O I] and O I abundances from our data (or that of Clegg *et al.* 1981). We emphasize that this result is not

due to different parameter values used here. Indeed, the slightly lower (on average) T_{eff} estimates we employ should result in slightly larger O I abundances and slightly lower [O I] abundances compared with NE92. We would then expect, all other things being equal, our $\Delta[\text{O}/\text{H}]$ values to be larger than those of NE92.

3.5.3 $\Delta[\text{O}/\text{H}]$ vs. $[\text{Fe}/\text{H}]$

NE92 find that their differences between the 6300 and 7774 Å O abundances increase with decreasing $[\text{Fe}/\text{H}]$. This is in the sense predicted for NLTE effects by Kiselman's (1991) calculations, but could also be due to inadequacies in model atmospheres (*e.g.*, the effects of inhomogeneities) which may increase at lower metallicities. This trend persists in our reanalysis of the NE92 data. A trend of $\Delta[\text{O}/\text{H}]$ in Figure 30a, which shows NE92's reanalyzed data, is evident. The ordinary correlation coefficient indicates the trend is real at the 98.8% confidence level. Figure 30b shows the remainder of the data analyzed in this chapter. No correlation is obvious (the statistical significance is at $< 50\%$ confidence level), but the data do not reach to as low $[\text{Fe}/\text{H}]$ as the NE92 data. Figure 30c shows the data sets plotted together.

3.5.4 Is There a 7774–6300 Å Discrepancy?

The analysis of our own data as well as that of the Clegg *et al.* (1981) suggests that there are no significant differences between 6300 Å and 7774 Å abundances for stars having $T_{\text{eff}} \lesssim 6200 - 6300$ K. However, the metallicity range of this sample is limited and no stars have $[\text{Fe}/\text{H}] < -0.5$. Is there evidence for a discrepancy at lower metallicities? If the revised T_{eff} estimates of chapter 2 are used, the answer appears to be “no”. The mean [O/Fe] ratio found from the 7774 Å triplet in metal-poor dwarfs in chapter 2 is the same, to within the uncertainties, as the mean [O/Fe] ratio of metal-poor field giant sample shown in figure 2 of Bessell, Sutherland & Ruan (1991) (see chapter 4 or King 1993b). Nor does the mean [O/Fe] found in chapter 2 differ from that of the two dwarfs with $[\text{Fe}/\text{H}] < -1.0$ analyzed by Spite & Spite (1991), who used the 6300 Å [O I] line. These findings are consistent with the lack of significant NLTE effects on the 7774 Å abundances as found by the calculations of Tomkin *et al.* (1992). However, the uncertainties certainly do not allow us to rule out the presence of small (~ 0.05 dex, say) effects due to NLTE departures or thermal inhomogeneities.

We cannot account for the differences between NE92's data and the other data analyzed here. It most likely results from differences in the model atmospheres utilized in the analysis. In view of the theoretical results of Tomkin *et al.* (1992) and Abia & Rebolo (1989) concerning NLTE effects, can

NE92’s results be understood? We argue that they can. If their larger values of $\Delta[\text{O}/\text{H}]$ are correct, then the trend with decreasing $[\text{Fe}/\text{H}]$ may be due to the effects of atmospheric inhomogeneities rather than NLTE effects. The possibility of such inhomogeneities has been discussed before by several authors including NE92. We repeat their call for the hydrodynamical calculations of, *e.g.*, Nordlund & Dravins (1990) to be extended to metal-deficient stars. However we note the recent pseudo-two-stream calculations of Tomkin *et al.* (1992). For the case having hot and cold streams each 400 K different from the actual effective T_{eff} , they find that the 7774 Å O I abundance would be reduced by 0.09 dex for HD 103095. Even more interesting is the result that the abundance derived from the 6300 Å [O I] line would *increase* by 0.16 dex for the same star. Therefore, without more detailed calculations for guidance, it must be stated that the commonly held belief that atmospheric inhomogeneities only affect the permitted lines and make them useless as abundance indicators may not be true.

The most important conclusion concerning the alleged 7774–6300 Å discrepancy is the potential existence of substantial systematic uncertainties in the abundances derived from *both* lines. The uncertainty in the equivalent width of the solar 6300 Å [O I] line (the value used by Lambert 1978 *versus* the value derived from the KFBT solar atlas using the latter’s continuum level) leads to differences in derived O abundances of $\gtrsim 0.2$ dex. Systematic effects on the O abundance derived from the 6300 Å [O I] lines due to atmospheric inhomogeneities may be comparable or larger than those on abundances derived from the permitted triplet. Uncertainties in the model atmospheres apparently can lead to systematic effects of at least ~ 0.1 dex in O abundances derived from the 7774 Å triplet. There also possibly exist systematic uncertainties in the metal-deficient dwarf T_{eff} scale (chapter 2, King 1993a). Additionally, uncertainty in the *gf* values of the 7774 Å triplet apparently exist. The values adopted here are some 0.10 – 0.11 dex larger than those of NE92. The new theoretical values of Biemont *et al.* (1991) are ~ 0.15 dex larger than those used by NE92. These systematic effects may either mask or exaggerate alleged NLTE or inhomogeneity effects on abundances derived from the 6300 and 7774 Å lines.

We conclude by stressing three points: (1) While the 6300 Å transition is thought to be free of the effects of NLTE departures and inhomogenieties (which may not be true given the two-stream results noted above), it cannot be said conclusively that it yields “better” O abundances. There are potential systematic effects in 6300 Å abundances due to uncertainty in the solar equivalent width which make it unclear to us if abundances derived from this line are any more accurate than those derived from the 7774 Å triplet. (2) Our data indicates that the 7774 Å lines seem to provide just as accurate and precise abundances as the 6300 Å line for the more metal-rich stars.

Hence, for determining the O abundance of open clusters we will retain their use. Even if NLTE or inhomogeneity effects are present in the 7774 abundances (which we do not regard to be the case for metal-rich stars— at the least it has not been conclusively demonstrated), our analysis (which uses the Kurucz models, uses our adopted gf values, and carefully normalizes the stellar abundances to the solar abundance derived for each line— as opposed to adopting a single solar abundance) seems to remove their effects quite successfully. (3) Removing potential systematic effects in O abundances derived from the 6300 Å and 7774 Å lines may prove intractable. Hence, more attention should be paid to other transitions which may yield more reliable O abundances free of these systematic effects. The new generation of very high UV QE CCD detectors on the next generation of large telescopes should enable O abundances to be determined in even the most metal-poor dwarfs from UV OH lines. Additionally, the new cryogenic echelle spectrographs on IR-optimized telescopes make possible observations of infrared OH lines. We eagerly await the results of others who are pursuing these programs and a comparison of these UV and IR abundances with those from optical lines.

4. Summary and Conclusions

We have investigated systematic effects on the O abundances derived from the 6300 Å [O I] line and 7774 Å O I triplet in solar-type stars. A summary of our findings is as follows:

- (1) Our 7774 Å Solar data, though apparently affected by the presence of scattered light in the spectrograph, confirm the flatness of the equivalent width– μ relation of Altrock (1968) as compared to that predicted by LTE models. The NLTE models of Tomkin *et al.* (1992) seem to fit the variation of equivalent width with μ better than LTE models even though the former are steeper than the observations at small μ .
- (2) The equivalent widths of the 7774 Å triplet measured from *flux* spectra are identical to those of both observed and model LTE *intensity* spectra at $\mu = 0.61 - 0.66$. This is result expected by employing the Eddington-Barbier relation. Since we know of no reason why these classic assumptions should break down in more metal-poor stars for the 7774 Å triplet, it seems reasonable to presume that accurate O abundances can be derived using an LTE analysis of the O I triplet for subdwarfs, all other things being equal (*i.e.*, assuming that other effects, such as thermal inhomogeneities or NLTE departures, do not become prevalent at lower metallicities).
- (3) There is good agreement between recent measurements of the Solar 7774 Å equivalent widths.

However, this does not appear to be the case for the 6300 Å [O I] line. Values range from 4.55 mÅ used by Lambert (1978) to ~ 7.9 mÅ, which we measure from the KFBT atlas using their continuum setting. These differences translate into abundance differences of $\gtrsim 0.2$ dex. Deriving Solar abundances or gf values from different data (*i.e.*, atlases) and/or using Solar continuum normalizations inconsistent with stellar data, may exacerbate the problem by potentially introducing additional systematic effects. These problems may be alleviated in the future by a) obtaining solar spectra from the bright sky or moon using the same instrumentation as for stellar data and b) using a stellar continuum normalization consistent with the Solar one. Given a) the uncertainty in the Solar 6300 Å equivalent width b) the systematic effects potentially introduced by using inhomogeneous Solar data and continuum normalizations, and c) the inability of Bessell *et al.* (1991) to reproduce others' 6300 Å abundances in their reanalysis, we (contrary to others) do *not* think it can be claimed *a priori* that the 6300 Å line provides more reliable O abundances than the 7774 Å triplet. While the 6300 Å line may be completely free of the effects of NLTE departures, it may be subject to systematic effects just as large as those that have been claimed to affect 7774 Å abundances.

(4) Comparison of O abundances derived from the 6300 and 7774 Å lines in a sample of relatively metal-rich F and G dwarfs reveals no systematic differences for $T_{\text{eff}} \lesssim 6200 - 6300$ K. At higher T_{eff} , the 7774 Å abundances are significantly larger. This likely reflects the onset of well-established and significant NLTE effects in early-type stars. *If* any systematic differences exist between 6300 Å and 7774 Å abundances in the stars we studied, our analysis (which is specific to our particular choice of oscillator strengths, software, model atmospheres, and procedure for normalization to the Solar abundance) seems to have removed them.

(5) The last result differs from that of NE92, who find significant differences between the 6300 Å and 7774 Å abundances. These differences persist in our reanalysis of their data. Likely causes are differences in the model atmospheres used (note that our “reanalysis” of their 7774 Å data is *not* independent of their model atmospheres since NE92 do not provide their equivalent widths) and perhaps also a) their procedure for normalizing abundances to Solar and b) their procedure for deriving 7774 Å gf values.

(6) The fact that our data (and that of CLT) and analysis reveals no systematic differences between the 6300 Å and 7774 Å abundances (for $T \lesssim 6200 - 6300$ K, of course) is consistent with the calculations of Tomkin *et al.* (1992), who find that NLTE effects on the 7774 Å abundances are not significant (again for $T_{\text{eff}} \lesssim 6200 - 6300$ K). If the hotter T_{eff} values advocated for metal-poor dwarfs in chapter 2 are correct, then there also appears to be no significant difference between O abundances in these

dwarfs as derived from the 7774 Å triplet and the abundances in metal-poor giants derived from the 6300 Å line. This finding would seem to preclude the existence of other substantial systematic effects in the 7774 Å abundances due to, *e.g.*, thermal inhomogeneities. However, we can not exclude the existence of rather small ($\lesssim 0.05$ dex) systematic effects. We believe that it remains to be conclusively demonstrated that the 7774 Å triplet provides less reliable abundances (again, for $T_{\text{eff}} \lesssim 6200 - 6300$ K) than the 6300 Å [O I] line in either metal-rich or metal-poor solar-type stars.

Table 9. Solar O I Triplet Equivalent Widths

μ $\cos \theta$	7772 Å mÅ	7774 Å mÅ	7775 Å mÅ
1.00	74.5	66.9	56.5
0.98	73.5	63.9	53.3
0.94	70.8	62.5	50.9
0.88	68.7	59.8	50.2
0.80	67.7	56.7	46.8
0.74	65.3	54.9	44.0
0.68	62.6	54.3	42.8
0.50	58.0	48.2	38.1
0.30	54.9	44.5	36.1
0.23	52.4	43.8	34.0

Table 10. Comparison of Solar 7774 Å O I Equivalent Widths

Source	7772 Å ($\mu = 1$) mÅ	7774 Å mÅ	7775 Å mÅ	7772 Å (Integrated) mÅ	7774 Å mÅ	7775 Å mÅ
Altrock	84.8	72.9	59.4	—	—	—
This Paper [†]	78.4	70.4	59.5	—	—	—
Utrecht Atlas	75	66	50	—	—	—
Liege Atlas	80.4	69.1	54.6	—	—	—
Edvardsson <i>et al.</i>	—	—	—	69.2	60.3	47.3
KFBT	—	—	—	77.1	66.6	52.0
AMB	—	—	—	71.5	62.0	45.7
JRK	—	—	—	72.0	59.0	45.3

[†]Corrected for 5.0% scattered light.

Table 11. Stellar Parameters

Star	T_{eff} K	$\sigma(T_{\text{eff}})$ K	$\log g$	$\sigma(\log g)$	[Fe/H]	ξ km/s
HR 33	6156	45	4.12	0.16	-0.40	1.85
HR 219	5900	80	4.41	0.16	-0.27	1.30
HR 483	5883	30	4.25	0.16	+0.02	1.48
HR 573	6176	78	4.28	0.16	-0.34	1.70
HR 781	6409	23	4.22	0.15	-0.21	1.95
HR 870	6454	45	4.10	0.15	+0.20	2.15
HR 996	5673	79	4.27	0.21	+0.03	1.30
HR 1010	5821	57	4.33	0.17	-0.21	1.35
HR 1101	5983	32	4.16	0.15	-0.10	1.70
HR 1294	5745	57	4.17	0.15	-0.18	1.50
HR 1536	5926	66	3.99	0.15	+0.14	1.85
HR 1673	6386	53	4.10	0.17	-0.30	2.10
HR 1729	5852	32	4.11	0.15	+0.03	1.65
HR 2354	5787	87	4.11	0.16	+0.13	1.60
HR 2493	6010	52	4.25	0.15	-0.37	1.60
HR 2548	6369	32	4.12	0.15	-0.20	2.05
HR 2883	5905	46	4.18	0.15	-0.77	1.60
HR 2943	6585	26	4.07	0.16	-0.02	2.30
HR 3018	5806	25	4.38	0.17	-0.84	1.25
HR 3775	6344	30	4.11	0.15	-0.15	2.05
HR 4529	6074	45	4.04	0.15	+0.16	1.90
HR 4533	6304	24	4.18	0.15	+0.20	1.90
HR 4540	6149	20	4.16	0.16	+0.13	1.80
HR 4734	5710	41	4.17	0.18	+0.10	1.45
HR 4903	5915	46	4.00	0.17	+0.24	1.80
HR 4983	6009	41	4.31	0.17	+0.06	1.50
HR 5868	5915	27	4.18	0.16	-0.03	1.60
HR 5933	6276	42	4.25	0.15	-0.14	1.80
HR 6212	5754	79	3.95	0.15	+0.07	1.75
HR 6467	6405	7	4.20	0.16	-0.42	1.95
HR 6541	6228	42	3.96	0.16	-0.21	2.10
HR 7322	6313	81	4.16	0.15	-0.28	1.95
HR 7462	5237	47	4.65	0.35	-0.24	0.80
HR 7503	5790	77	4.10	0.15	+0.08	1.60
HR 7504	5719	63	4.09	0.16	+0.03	1.55
HR 7534	6345	45	4.22	0.15	-0.13	1.90
HR 7672	5880	85	4.34	0.19	-0.11	1.35
HR 8077	6116	16	4.08	0.16	-0.09	1.90
HR 8181	6134	67	4.35	0.16	-0.71	1.55
HR 8314	5926	69	4.30	0.15	+0.03	1.45
HR 8805	6469	32	4.21	0.15	-0.14	2.00
HR 8825	6366	69	4.15	0.16	-0.20	2.00
HR 8969	6157	49	4.17	0.15	-0.21	1.80

Table 11. (Continued) Stellar Parameters

Star	T_{eff} K	$\sigma(T_{\text{eff}})$ K	$\log g$	$\sigma(\log g)$	[Fe/H]	ξ km/s
HD 22879	5859	74	4.29	0.15	-0.84	1.40
HD 51929	5853	70	4.26	0.15	-0.72	1.45
HD 78558	5707	31	4.23	0.17	-0.40	1.35
HD 89707	5966	34	4.42	0.15	-0.42	1.35
HD 98553	5898	59	4.36	0.15	-0.43	1.35
HD 130551	6251	35	4.26	0.15	-0.62	1.75
Sun	5770	15	4.44	0.1	+0.00	1.15

Table 12. Equivalent Widths and Abundances

Star	EW(6300) mÅ	σ mÅ	Ref	EW(7774) mÅ	σ mÅ	Ref	log ϵ (O) 6300Å	σ dex	log ϵ (O) 7774Å	σ dex
HR 33	5.9	1.8	JRK	242.2	7	JRK	8.76	0.18	8.61	0.07
HR 219	4.4	0.5	CLT	185.0	13	CLT	8.74	0.09	8.77	0.11
	4.6	-	SLW	178.0	-	SLW	-	-	-	-
HR 483	6.2	0.8	CLT	192.2	20	CLT	8.93	0.07	8.76	0.12
	6.8	1.2	JRK	213.8	7	JRK	-	-	-	-
HR 573	3.55	0.15	NE	-	-	NE	8.60	0.08	8.63	-
HR 781	6.5	2.5	JRK	341.0	8	JRK	8.94	0.24	8.97	0.05
HR 870	9.2	2.8	JRK	442.8	7	JRK	9.20	0.13	9.28	0.06
HR 996	7.4	1.7	JRK	174.5	6	JRK	8.97	0.15	8.88	0.11
HR 1010	3.45	0.6	NE	-	-	NE	8.62	0.12	8.65	-
HR 1101	5.25	0.5	NE	-	-	NE	8.81	0.08	8.90	-
HR 1294	5.1	0.4	NE	-	-	NE	8.74	0.08	8.72	-
HR 1536	7.55	0.75	NE	-	-	NE	8.97	0.08	8.91	-
HR 1673	4.6	0.5	GO	320.0	14	CLT	8.72	0.11	8.86	0.08
HR 1729	7.6	0.8	CLT	219.0	14	CLT	8.97	0.09	8.87	0.10
	-	-	-	228.0	-	SLW	-	-	-	-
HR 2354	7.85	0.45	NE	-	-	NE	8.97	0.07	8.94	-
HR 2493	3.65	0.35	NE	-	-	NE	8.55	0.09	8.52	-
HR 2548	3.5	0.1	NE	-	-	-	8.63	0.06	-	-
HR 2883	3.9	0.4	NE	-	-	NE	8.46	0.08	8.56	-
HR 2943	3.7	0.3	NE	-	-	NE	8.74	0.08	9.07	-
HR 3018	3.35	0.15	NE	-	-	NE	8.42	0.07	8.62	-
HR 3775	4.7	0.5	CLT	342.0	18	CLT	8.76	0.08	8.96	0.08
HR 4529	9.5	0.2	NE	-	-	-	9.13	0.07	-	-
HR 4533	6.85	0.15	NE	-	-	-	9.08	0.06	-	-
HR 4540	6.6	0.7	CLT	309.0	12	CLT	9.01	0.08	9.05	0.07
	6.4	0.2	NE	-	-	NE	9.01	0.06	8.96	-
HR 4734	8.37	0.85	NE	-	-	NE	9.00	0.09	8.90	-
HR 4903	10.85	0.15	NE	-	-	NE	9.13	0.07	9.16	-
HR 4983	5.7	0.6	CLT	229.0	10	CLT	8.97	0.08	8.86	0.07
HR 5868	8.5	0.8	AMB	226.4	12	AMB	9.06	0.05	8.88	0.08

Table 12. (Continued) Equivalent Widths and Abundances

Star	EW(6300) mÅ	σ mÅ	Ref	EW(7774) mÅ	σ mÅ	Ref	log $\epsilon(\text{O})$ 6300Å	σ dex	log $\epsilon(\text{O})$ 7774Å	σ dex
HR 5933	6.0	0.9	CLT	275.0	15	CLT	8.92	0.10	8.81	0.08
HR 6212	12.4	1.2	AMB	272.0	13	AMB	9.09	0.08	9.15	0.09
HR 6467	2.7	0.5	AMB	298.3	10	AMB	8.51	0.10	8.79	0.06
	2.7	0.8	JRK	304.6	7	JRK	-	-	-	-
HR 6541	6.0	0.6	AMB	332.0	17	AMB	8.76	0.07	8.97	0.09
HR 7322	6.3	0.6	AMB	319.3	16	AMB	8.86	0.08	8.94	0.10
HR 7462	3.6	0.6	CLT	69.0	4	CLT	8.62	0.16	8.68	0.13
HR 7503	5.4	0.5	AMB	209.8	7	AMB	8.79	0.08	8.87	0.09
	5.8	0.9	JRK	197.8	4	JRK	-	-	-	-
HR 7504	6.7	0.6	AMB	186.9	10	AMB	8.87	0.08	8.82	0.10
	6.8	1.1	JRK	-	-	-	-	-	-	-
HR 7534	4.1	1.6	JRK	324.5	7	JRK	8.77	0.23	8.95	0.05
HR 7672	6.8	1.3	JRK	219.4	6	JRK	8.97	0.15	8.90	0.08
HR 8077	8.5	1.0	JRK	313.8	5	JRK	9.01	0.09	9.03	0.04
HR 8181	3.0	0.6	BE	-	-	E93	8.49	0.13	8.39	-
HR 8314	5.7	0.8	AMB	220.1	11	AMB	8.92	0.10	8.89	0.09
HR 8805	4.3	1.1	JRK	379.0	9	JRK	8.80	0.23	9.08	0.06
HR 8825	2.7	0.4	AMB	343.7	25	AMB	8.53	0.10	8.99	0.12
HR 8969	6.2	0.6	AMB	290.7	10	AMB	8.85	0.09	8.96	0.05
	-	-	-	284.3	4	JRK	-	-	-	-
HD 22879	2.9	0.3	NE	-	-	NE	8.34	0.08	8.51	-
HD 51929	3.95	0.1	NE	-	-	NE	8.49	0.07	8.61	-
HD 78558	6.5	0.3	NE	-	-	NE	8.78	0.06	8.86	-
HD 89707	3.3	0.5	NE	-	-	NE	8.56	0.10	8.61	-
HD 98553	3.1	0.3	NE	-	-	NE	8.51	0.08	8.44	-
HD 130551	2.8	0.5	NE	-	-	NE	8.47	0.11	8.53	-
Sun	5.94	0.3	K84	179.2	4	AMB	8.95	0.02	8.89	0.02

References: 1) AMB=CFHT data from this chapter 2) BE=Barbuy & Erdelyi-Mendes (1989)

3) CLT=Clegg *et al.* (1981) 4) E93=Edvardsson *et al.* (1993) 5) GO=Gratton & Ortolani (1986)

6) JRK=UH 2.2-m data from this chapter 7) K84=Kurucz *et al.* (1984) atlas 8) NE92=Nissen & Edvardsson (1992)

Figure 22. Solar spectra of the 7774 Å region we acquired at Mees Solar Observatory. The spectra were taken at different values of μ and have been offset by an arbitrary constant for clarity. The O I lines at 7772, 7774, and 7775 Å are clear. The deep feature at 7781 Å is an Fe I line.

Figure 23. a) The equivalent width of the 7772 Å O I line as measured by Altrock (1968, open symbols) and us (filled symbols) is plotted against μ ($= \cos \theta$, where θ is the apparent angle measured from disk center to the limb. The solid line is the LTE model prediction assuming a near-meteoritic O abundance. b) Same as (a) except for the 7774 Å O I line. c) Same as (a) except for the 7775 Å line. d) The data of Altrock (1968) alone (symbols) for all three lines (from top to bottom: 7772, 7774, and 7775 Å) is plotted versus μ with our LTE model results for each line (solid lines).

Figure 24. The data of Altrock (1968, symbols) for each line in the O I triplet are plotted against μ . The solid lines are the predictions of a NLTE calculation by Tomkin *et al.* (1992).

Figure 25. A 7774 Å region spectrum of the star HR 8969 taken at the UH 2.2-m. The O I lines at 7772, 7774, and 7775 Å are marked by the vertical lines.

Figure 26. The difference in the $\log g$ values of E93 and those predicted from Stromgren colors of the new Kurucz (1992) models are plotted against the Kurucz model values. The solid line is an ordinary least squares regression to the data points and is used to place the Kurucz model values on the scale of E93.

Figure 27. The $[O/H]$ ratio as derived from the 6300 Å $[O\ I]$ line is plotted against the $[O/H]$ ratio derived from the 7774 Å $O\ I$ lines. The open squares are the reanalyzed data of NE92. The filled squares represent all other data used here. The dashed line represents the relation expected if the two $[O/H]$ ratios are equal. It can be seen that the 7774 Å $O\ I$ lines appear to yield larger O abundances.

Figure 28. a) $\Delta[\text{O}/\text{H}]$, defined as $[\text{O}/\text{H}](7774 \text{ \AA}) - [\text{O}/\text{H}](6300 \text{ \AA})$, is plotted against our effective temperature for the stars of NE92 which are reanalyzed here. b) Same as (a) except showing all the data in Table 4 other than NE92's. An increase in $\Delta[\text{O}/\text{H}]$ for $T_{\text{eff}} \gtrsim 6200$ K is seen. c) The data in (a) and (b) are plotted together.

Figure 29. Same as Figure 28a except the original abundances and T_{eff} values of NE92 are used. The significantly positive value of $\Delta[\text{O}/\text{H}]$ seen in Figure 28a persists.

Figure 30. a) $\Delta[O/H]$ is plotted *versus* $[Fe/H]$ for the reanalyzed NE92 data. A significant trend of increasing $\Delta[O/H]$ with declining $[Fe/H]$ is seen. b) Same as (a) except showing all the data in Table 4 other than NE92's. No trend with declining $[Fe/H]$ is seen, but the range of $[Fe/H]$ is limited. c) The data in (a) and (b) are plotted together.

Chapter 6. Oxygen Abundances in Open Clusters

Abstract

Oxygen abundances are determined for dwarfs or subgiants in the Pleiades, Coma, UMa, Praesepe, Hyades, NGC 752, M 67, and NGC 188 open clusters (or moving groups). The mean abundances for the Hyades and the UMa group are in poor agreement with the recent determinations of Garcia Lopez *et al.* (1993). We discuss why we believe our abundances to be more accurate. We find a very clear global relation between cluster age and O abundance over the past 7 Gyr. This behavior is in stark contrast to the lack of any correlation between Fe abundance and cluster age. This lack of correlation and the presence of significant scatter demonstrated by field stars in the $[\text{Fe}/\text{H}]$ vs. age plane may be consistent with CO-He white dwarf systems comprising the majority of Type Ia supernova systems. If this is correct, then the duration of halo formation may be several Gyr— long enough to accommodate age spreads among globular clusters and halo field stars. It is concluded that despite possible *a priori* objections about using O as a chronometer, it may be superior to Fe in this regard. Further determinations of O abundances in open clusters are urged in order to confirm the results presented here and to investigate abundance gradients in the disk. The O abundances in the younger clusters are super-solar and are significantly larger than those determined in H II regions, planetary nebulae, and supergiants. It is suggested, and supported with observational evidence, that this may be due to: 1) a lack of consideration of additional mixing processes in stellar models such that the observed O abundances in planetary nebulae and supergiants do not reflect their initial abundances and 2) the possibility that a substantial fraction of O in gaseous nebulae is locked up in dust grains. Keeping in mind the possible incompleteness in stellar models, we reinterpret the existing planetary nebulae and supergiant abundances as being in good agreement with our young open cluster results. Mixing processes currently not considered in stellar models may also mean that O abundances derived in halo giants have been affected. Again, keeping this in mind, we reinterpret existing red giant O data as indicating $[\text{O}/\text{Fe}] \sim +0.5$ in the halo. We suggest that investigation of models, such as the Yale rotational ones, is urgently needed in both metal-rich and metal-poor stars having a variety of masses to shed further light on these subjects. If the proposed resolution to the present-day O abundance discrepancy is correct, then inferences made about extragalactic O abundances or gradients in the Galactic O abundance may be incorrect.

1. Introduction

The Galaxy’s population of open clusters has long served as a probe of the structure, dynamics, and chemical evolution¹¹ of the Galactic disk (Friel & Janes 1993). The major advantage in studying the history and properties of the disk using fainter cluster stars instead of brighter disk field stars is the belief that the former can be dated with much more reliability than the latter. Thus, in the investigation of such issues as the disk’s age-metallicity relation(s), one hopes that the clusters’ color-magnitude diagrams provide a more accurate (and precise) chronometer. Indeed, the age determination of field stars may play an important role in differences between various age-metallicity relations derived for the disk (*e.g.*, the results of Twarog 1980 *versus* those of Barry 1988). The main drawback of clusters is that the older ones lie at large distances from the Sun (whereas field star samples are usually restricted to the so-called Solar “neighborhood”). This presents an admitted problem if parameters such as infall rates, the IMF, SFR, and mixing times are all functions (even if simple ones) of Galactic radius, azimuth, and/or height above the plane (Sandage 1988). One must be prepared to deconvolve all these possible effects, a difficult task at best and one which requires a large sample size, if one is to better understand the Galactic disk.

For some time, abundances for the more distant clusters have been determined from photometry of cluster giants in the DDO and UBV system (Janes 1979) and the Washington photometric system (Geisler, Claria & Minniti 1992). Intermediate resolution spectra have provided a wealth of information on cluster kinematics and abundances (Friel & Janes 1993). With the availability of high-efficiency, linear detectors, it has now become possible to determine precise abundances for cluster dwarfs (Cayrel, Cayrel de Strobel & Campbell 1985; Boesgaard 1989; Boesgaard & Friel 1990; Friel & Boesgaard 1990) from high S/N, high resolution spectroscopy— even in more distant clusters such as M67 (Friel & Boesgaard 1992). Using data taken with early generation CCDs (whose QEs pale in comparison to newer generation devices), it is possible to accurately determine effective temperatures and gravities (and abundances) *spectroscopically* for the important old and more distant clusters such as NGC 188 (Hobbs, Thorburn & Rodriguez-Bell 1990), M67 (Hobbs & Thorburn 1991), and NGC 752 (Hobbs & Thorburn 1992). The renewed interest in spectroscopy of open cluster members is now partially fueled by the availability of multi-object spectrographs. It is now a rather trivial matter to obtain 50-100 spectra on an individual CCD frame per exposure.

¹¹ It must also be mentioned that the clusters provide extremely important benchmarks in the investigation of stellar evolution and stellar atmosphere/interior physics by providing “snapshots” of powerful tracers such as Li and Be abundances and $^{12}\text{C}/^{13}\text{C}$ ratios.

The element traditionally used in the study of Galactic chemical evolution is Fe, due to its relative ease of study in stellar spectra. Wheeler, Sneden & Truran (1989) have suggested that Fe may not be a useful chronometer. They note that the potential problems with utilizing iron abundances are: a) the fact that there are two distinct sources for Fe production— one from high mass progenitors and another from low mass progenitors; hence there is a blurring of time scales. b) Theoretical yields for iron are somewhat uncertain due to lack of precise estimates of how much ejecta is blown away from the Galaxy or how much falls back onto a compact remnant. As Wheeler *et al.* (1989) state, O may avoid these problems as it likely has one predominant source and its theoretical yield can be much more securely computed. We reiterate their caution, though, that the use of O as a chronometer in the disk may have different and significant drawbacks. Of particular note for the open cluster population is that the history of the O abundance in the disk can be seriously affected by spatial and/or temporal variations in the IMF and SFR (*i.e.*, starbursts). However, to avoid studying the O abundances of the Galaxy’s open clusters due to these uncertainties (especially when the uncertainties using Fe may be just as great) would be short-sighted.

Unfortunately, determinations of O abundances in open clusters have been few in number compared to those in field stars and the determination of other elemental abundances in clusters. The only works that we are aware of are those of Garcia Lopez *et al.* (1993), who determined O abundances in Hyades and UMa group F dwarfs, and Brown (1985) and Cohen (1980), who determined O abundances in M 67 giants. It is particularly surprising that solar-age and younger clusters have not had their O abundance studied further since it is well known that observations of B stars, H II regions, supergiants, and planetary nebulae yield O abundances much lower (by ~ 0.2 dex) than those observed in the Sun. This paper is an attempt to remedy the lack of O abundance determinations in open cluster dwarfs. Here we present O abundances based on moderate resolution, high S/N observations of the 7774 Å O I triplet in relatively cool dwarf members of the Pleiades, Coma, the UMa moving group, Praesepe, the Hyades, NGC 752, M 67 (NGC 2682), and NGC 188.

2. Observations and Data Reduction

2.1 UH 2.2-m Observations

Observations of the 7774 Å O I triplet in relatively cool dwarfs (roughly 5800-6200 K) in the Pleiades, Coma, UMa, Praesepe, and Hyades clusters (or moving group in the case of UMa) were obtained with the University of Hawai’i 2.2-m meter telescope on Mauna Kea during several observing runs during 1991-1993. The data were obtained with the Faint Object Spectrograph

which employed a 400 mm camera, 1200 l/mm grating blazed in the first order at 6300 Å, a 0.7" slit, and either a Tektronix 1024×1024 or 2048×2048 CCD with 24 μm pixels. This instrumental configuration yielded a dispersion of ~0.50 Å/pixel. The resolution, measured from the FWHM of Fe-Ar comparison lamp lines, was 1.2 – 1.3 Å. Flat-field frames consisted of ~2 minute exposures of a Tungsten lamp at the top of the telescope dome. Object exposure times ranged from 5 to 30 minutes and the final S/N (per pixel, in the continuum region near the 7774 Å O I triplet) ranged from 200 to 500.

2.2 KPNO 4-m Observations

Observations of dwarfs (and, technically, little-evolved subgiants) in NGC 188, NGC 752, and M 67 were carried during two nights in October 1992 at the KPNO 4-m. The data were obtained with the Hydra multi-object fiber feed device, red fiber cables, the mounted bench spectrograph, the Simmons camera, the 1200 l/mm grating #380, an RG 610 filter (to block light from higher orders), and the T2KB (2048×2048, 24 μm pixels) CCD. This configuration also yielded a dispersion of ~0.50 Å/pixel, but the resolution, 2.0 – 2.3 Å, was coarser than the UH 2.2-m observations. Sufficient calibration frames (bias, dome flats, and Th-Ar lamps) were obtained each night. Besides the regular beginning/ending of night dome flats, flat field exposures were also obtained for each specific fiber configuration. Total exposure times of 225, 122, and 342.5 minutes were achieved for the NGC 752, M 67, and NGC 188 fields. The final median S/N of the coadded, 1-d spectra was 142, 103, and 94 for NGC 752, M 67, and NGC 188 as determined from Poisson noise. The measured S/N is lower (~75), however, for M 67 and NGC 188 due to limitations in the sky subtraction. The realized S/N for the M 67 data is similar to that for NGC 188 due to a smaller number of fibers used in the sky subtraction. A lower number was used as an expedient in configuring the Hydra fibers as quickly as possible in order to obtain as much integration time on M 67 as possible before sunrise.

Astrometry for the NGC 752 field was performed using the measurements in Heinemann (1926). SAO coordinates for 8 of the brighter stars in the Heinemann (1926) list were used to transform his plate coordinates to α and δ . The resulting positions proved to be very accurate as only a couple of the Hydra fibers required slight “tweaking” after setting upon the field acquisition stars. For M 67, we used the positions given by Girard *et al.* (1989). This astrometry was excellent and no tweaking was required for any of the fibers we checked. Dr. K. Janes kindly provided a V band CCD image of NGC 188 which we used to determine stellar positions. Our “plate” solutions were not as accurate as possible (probably due to the small number of positional standards used) and the

astrometry was not as good as for the other clusters and about 25 fibers had to be tweaked before exposing. However, it is of some importance to nonetheless note that, after taking into account our marginal astrometric solutions, Hydra performed perfectly fine at the extreme declination of NGC 188 ($\delta = +85$).

2.3 Reductions

Reductions of the UH 2.2-m data followed standard procedures within the IRAF software package. This included removal of the dc-offset of the detector by using the median or mode of the counts in the overscan region. No additional bias pattern removal was attempted as it is known to be variable (in magnitude and structure) for the particular CCDs used. However, the additional noise that the variability adds is dwarfed by ordinary photon noise. Flat-fielding was accomplished by first averaging (after 2.5σ clipping [from the median] deviant pixels) usually 4 – 8 frames of Tungsten lamp exposures. These frames were then fit with a low to moderate order spline to remove large scale structure and achieve proper normalization. The flat fields were then divided into the overscan-corrected object frames to remove pixel-to-pixel variations of the detector. Tests indicated that the flat-fielding was usually good to within 0.4 – 0.5% per *unsummed* pixel. Residual background (sky, the negligible dark current, *etc.*) was removed by fitting (in the direction parallel to the dispersion) a low order polynomial to the average of 25-30 pixels displaced from the wings of the object spectrum’s spatial profile. In almost all cases, OH sky emission posed no problem to accurately measuring equivalent widths. In the few cases where the lines might have posed a problem to continuum placement, they were successfully removed using the “background” package in IRAF. The spectra were traced and summed using the usual “apextract” procedure in IRAF. Wavelength calibration was performed by fitting low to moderate order polynomials to 5 – 9 lines from the Fe-Ar exposures. The typical residuals of these fits were typically $\sim 0.006 \text{ \AA}$.

The Hydra runs were performed as detailed in the NOAO publication “Guide to the HYDRA Reduction Task DOHYDRA” by Francisco Valdes. The reductions began with subtracting the median of the count distribution in the overscan region from all frames. A nightly bias frame was constructed after averaging (after 2.5σ clipping outlying pixels) four or five individual overscan-subtracted bias frames. The bias frames were subtracted from all flat-field, comparison, and object exposures. The remainder of the reductions were carried out using the “dohydra” package in IRAF. This package performs the following steps: defines the apertures; fits the fiber positions; extracts, normalizes, and throughput corrects the flat field spectra; determines dispersion solutions from the comparison lamp exposures; extracts, flat-fields, and dispersion corrects the object spectra; and sky

subtracts the object spectra.

3. Abundance Analysis

3.1 The Cluster Sample

The first column of Tables 13–20 list the cluster stars observed in the Pleiades, Coma, the UMa moving group, Praesepe, the Hyades, NGC 752, M 67, NGC 188. The Pleiades stars are designated with their Herzberg numbers (H II numbers, not H I). Hyades stars are designated by van Buren numbers. All the stars in the Pleiades and Hyades are thought to be assured members based on examination of the relevant literature (*e.g.*, Griffin *et al.* 1988 for the Hyades).

The Coma stars are listed according to their Trumpler numbers. No modern membership analyses of this cluster seems to have been carried out. The three Coma stars we deliberately observed for this program all have very similar O abundances and we consider them assured members of the cluster. The fourth Coma star, Tr 23 (G 59-5), was observed as part of another program and it was only realized later that it was a suggested member of the cluster. Its O abundance (see below) is clearly inconsistent with the other three, but its [O/Fe] ratio is that expected of a metal-poor ($[M/H] \sim -0.78$; Laird, Carney & Latham 1988) field star.

The Praesepe star designations are predominantly those of Klein-Wassink (1927). Alternate designations, given in the second column of Table 16, are those of Jones & Cudworth (1983), Jones & Stauffer (1991) (who provide a convenient table of multiple cross-identifications), and Artjukhina (1966, 1971). The third column of Table 16 gives membership probabilities for the Praesepe stars based on the proper motion study of Jones and Stauffer (1991); all but one of the stars in our Praesepe sample seem to be assured members.

The UMa group members are designated with their HR or HD numbers. The membership of objects in this stellar aggregate is a somewhat thorny issue. Indeed, the issue of the reality of moving groups is a mildly contentious one itself. We discuss the UMa stars on an individual basis in a later section when it becomes time to adopt representative O abundances for an individual cluster as a whole. For the time being, suffice it to say that we drew potential UMa members from the studies of Johnson & Soderblom (1987), Soderblom & Mayor (1993), Boesgaard, Budge & Burck (1988), and Garcia Lopez *et al.* (1993). Another candidate moving group is the Wolf 630 aggregate noted by Eggen (1969, 1971). Although no deliberate attempt was made to observe candidate member stars of the Wolf 630 group as part of this program, we did observe a few as part of other related programs. Additionally, O abundances have been derived for some Wolf 630 group candidate members by Edvardsson *et al.* (1993). We initially sought to include this important potential moving group (as

it is proposed to be of the old disk population). However, further examination of even the best candidate members revealed a marked heterogeneity of the O and Fe abundances, which leads us to question the reality of this group. In any case, this heterogeneity means this aggregate can not be used in the present study (where we assume the clusters can be assigned a unique value of $[\text{Fe}/\text{H}]$ and $[\text{O}/\text{H}]$ — or at least a single value with perhaps a small real spread). Therefore, we consider Wolf 630 no further.

NGC 752 objects are listed by their Heinemann (19??) numbers. Listed in the second and third columns of Table 18 are the membership probabilities according to the proper motion studies of Platais (1991) and Francic (1989). All the stars in our sample appear to be certain members based on these studies and the derived O abundances.

M 67 objects are tabulated by their designation in Sanders (1977). Columns two, three, and four of Table 19 give the membership probabilities from the proper motion study of Sanders (1977), the proper motion study of Girard *et al.* (1989), and the combined proper motion—radial velocity analysis of Girard *et al.* (1989). It seems that all but one of the stars is a certain M 67 member.

The NGC 188 objects are labelled using the designation in Sandage (1962). Column two of Table 20 lists the membership probabilities based on the proper motion study of Uggren, Mesrobian & Kerridge (1972). It seems likely that all are cluster members based on these probabilities.

3.2 Stellar Parameters

3.2.1 $[\text{Fe}/\text{H}]$

Values for the clusters' iron abundance, $[\text{Fe}/\text{H}]$, were drawn from the literature. The values for the Pleiades, Coma, the UMa group, Praesepe, and the Hyades are taken from various metallicity determinations in the literature— namely, those of Boesgaard (1989); Boesgaard & Friel (1990); Boesgaard, Budge & Burck (1988); Boesgaard & Budge (1988); and (for the Hyades) Cayrel *et al.* (1985). What has actually been done here for these clusters is to adopt a Hyades abundance based on the literature estimates and then use the *relative* abundances of the other clusters to assign values of $[\text{Fe}/\text{H}]$ to them (*i.e.*, we fix the metallicity scale by using the Hyades value as a zero point). The precision Hyades iron abundances derived by Boesgaard & Budge (1988), Boesgaard (1989), and Boesgaard & Friel (1990) are somewhat higher than those derived by Cayrel *et al.* (1985) (though it must be stated that, even given the small uncertainties, this difference is probably not significant) by ~ 0.03 dex. Cayrel *et al.* (1985) do note, though, that their abundance of $[\text{Fe}/\text{H}] = +0.12$ dex could be a slight underestimate. Therefore, based on this information, we take a value of $[\text{Fe}/\text{H}] = +0.14$

as our *initial* best estimate. When one compares the values of the T_{eff} adopted here (see below), it seems that this value would have to be lowered slightly. We therefore adopt a Hyades abundance of $[\text{Fe}/\text{H}] = +0.12$ dex. Though this slight adjustment is insignificant in its bearing on our results and the O abundances themselves, it does provide a small measure of self-consistency.

Iron abundances for the older clusters NGC 752, M 67, and NGC 188 were taken from the analyses of Hobbs & Thorburn (1992), Hobbs & Thorburn (1991), and Hobbs *et al.* (1990). We regard these as probably the best values for these clusters owing to their spectroscopic T_{eff} estimates and the iterative analysis technique. It should be noted that these metallicities may or may not be on the same scale as the abundances adopted for the other clusters so that there is no guarantee of self-consistency on this account. However, the uncertainties in these values are relatively small and, as is well known, the O abundance derived from the 7774 Å O I triplet is insensitive to the assumed metallicity. Therefore, it is not believed that the use of heterogeneous sources for the iron abundances affects our conclusions. The clusters and the final adopted $[\text{Fe}/\text{H}]$ values are listed in columns one and two of Table 21, which summarizes the gross properties of our sample. Our best assessment of the 1σ uncertainties (including systematic effects) are given in the third column.

3.2.2 Cluster Reddenings

The cluster reddenings are needed in order to derive temperatures from *intrinsic* photometric indices. For Coma, the UMa group, Praesepe, and the Hyades, we make the usual assumption of zero reddening. The Pleiades is known to have significant (though not necessarily substantial) reddening and, moreover, suffers from definite variable reddening. Estimates of $E(B - V)$ for individual Pleiads were taken from Breger (1986). Values of $E(b - y)$ were taken from Crawford & Perry (1976). For each Pleiad, a final reddening was adopted based on the average of the Johnson and Stromgren values (assuming $E(b - y) = 0.72 \times E(B - V)$).

Hobbs & Thorburn (1992) use their spectroscopically derived T_{eff} values to estimate the intrinsic colors of NGC 752 stars using the relations of Saxner & Hammarback (1985). Their comparison of the intrinsic and apparent colors yielded an average reddening of $E(B - V) = 0.04$ mag. This value is in excellent agreement with the $E(b - y)$ values derived by Crawford & Barnes (1970) and Twarog (1983). $E(B - V) = 0.04$ is thus adopted here.

Unfortunately, the reddening for M 67 is more uncertain. Hobbs & Thorburn's (1991) spectroscopic T_{eff} values indicate an average value of $E(B - V) = 0.065$. Montgomery, Marschall & Janes (1993) derive a value of $E(B - V) = 0.05$ from analysis of the M 67 color-color diagram.

Demarque, Green & Guenther (1992) utilize a value of $E(B - V) = 0.06$ in their M 67 age derivation via comparison with theoretical isochrones. Nissen, Twarog & Crawford (1987), however, derive a lower reddening, $E(b - y) = 0.023$ ($E(B - V) = 0.032$), based on $uvby\beta$ photometry. Here, we chose to adopt a value intermediate to that inferred by Nissen *et al.* (1987) and that inferred by Hobbs & Thorburn (1991). We assume $E(B - V) = 0.050$ ($E(b - y) = 0.036$). Note that our T_{eff} values will then not (*on average*) be expected to agree with those of Hobbs & Thorburn (ours should be, again on average, slightly cooler). Therefore, their Fe abundance for M 67, which we adopt here, may not be strictly consistent with our temperatures. However, the issue is clouded as the abundances are actually derived from Fe II lines (which are relatively insensitive to T_{eff}) and then an iterative analysis is employed to demand consistency between the Fe II and Fe I abundances. In any case, the difference between our reddening and the average reddening implied by Hobbs & Thorburn's (1991) temperatures would imply a ~ 50 K difference on average. This would affect abundances derived from the Fe I lines by only a few hundredths of a dex. As noted above, an Fe abundance difference of this size does not alter our conclusions and does not affect our derived O abundance. Therefore we shall not be concerned further with this minor self-inconsistency.

A similar uncertainty in reddening exists for NGC 188. Eggen & Sandage (1969) inferred a value of $E(B - V) = 0.09$. Hobbs *et al.* (1990) inferred an average value of $E(B - V) = 0.10$ from their spectroscopic temperatures. However, Twarog & Anthony-Twarog (1989) argue for a differential reddening between M 67 and NGC 188 of $\Delta E(B - V) = 0.09$ in order to remove perceived inconsistencies in these clusters' Li abundances. With our reddening value for M 67 this translates to $E(B - V) = 0.14$ for NGC 188. We adopt a somewhat intermediate value of $E(B - V) = 0.11$ for NGC 188. An uncertainty in reddening of ± 0.01 mag leads to an uncertainty in the T_{eff} of roughly ∓ 40 K. Fortunately, the β index is reddening independent and should yield accurate T_{eff} values even if assumed reddenings are incorrect.

3.2.3 Effective Temperatures

Our effective temperature determinations followed the careful procedure used by Boesgaard, Budge & Ramsay (1988). Once cluster reddenings were adopted (or, in the case of the Pleiades, the specific stellar reddening was adopted), effective temperatures were determined for all the cluster stars using the photometric relations of Saxner & Hammarback (1985). Values from the $(B - V)$ and $(b - y)$ colors and the β index were averaged to form final estimates. For the Pleiades, $(B - V)$ photometry was taken from Johnson & Mitchell (1958). Stromgren photometry was taken from

Crawford & Perry (1976). $(B - V)$ colors for the Coma cluster was taken from Johnson & Knuckles (1955) while Stromgren photometry was taken from Crawford & Barnes (1969a). $(B - V)$ photometry for the UMa group candidates was taken from the *Bright Star Catalog* (Hoffleit & Jaschek 1982) and Stromgren photometry was taken from the Hauck & Mermilliod (1990) catalog¹². For Praesepe, the $(B - V)$ photometry was taken from Johnson (1952); the Stromgren photometry was from Crawford & Barnes (1969b) and Reglero & Fabregat (1991). Hyades $(B - V)$ photometry was taken from Griffin *et al.* (1988) and Carney (1983); the Stromgren photometry has multiple sources which are given in the Hauck & Mermilliod (1990) catalog.

For NGC 752, $(B - V)$ values were taken from the photoelectric photometry of Johnson (1953) and/or Eggen (1963) or, as a last resort, the photographic photometry of Francic (1989). Unfortunately, Stromgren photometry for the cooler members analyzed here was not available. $(B - V)$ photometry for M 67 was taken from the CCD results of Montgomery *et al.* (1993) and/or the photoelectric data of Eggen & Sandage (1964); the Stromgren photometry is predominantly from Nissen *et al.* (1987). For NGC 188, the $(B - V)$ photometry utilized consisted of the photographic results of Sandage (1962), the photoelectric values of Eggen & Sandage (1969), and/or the CCD results of Caputo *et al.* (1990). These were supplemented by unpublished CCD photometry kindly provided by Dr. K. Janes. The latter was extremely valuable to help sort out discrepant values from the other studies. Unfortunately, no Stromgren photometry is available for this cluster.

Based on the scatter of the T_{eff} values derived from the $(B - V)$, $(b - y)$, and $H\beta$ indices, we estimate the typical uncertainty of our T_{eff} values to be ~ 60 K for the younger clusters. For the older clusters (NGC 752, M 67, and NGC 188) the uncertainty is probably larger due to the lack of Stromgren photometry and the uncertainty in reddening. We estimate the uncertainty in these T_{eff} values to be $100 - 110$ K. This leads to an uncertainty in the derived O abundances of $\sim 0.08 - 0.09$ dex. The final values of T_{eff} are given in columns two of Tables 13, 14, 15, 17; column three of Table 20; column four of Tables 16 and 18; and column five of Table 19.

3.2.4 Gravities

Stellar gravities ($\log g$) were determined using Stromgren photometry in the same manner as in chapter 5. Stromgren photometry, taken from the sources listed above, was used to place the

¹² Actually, all the Stromgren photometry for all the clusters was taken from the Hauck & Mermilliod (1990) catalog. For the sake of completeness, however, we provide the *main* sources of this compilation above.

stars in the c_1 vs. $(b - y)$ plane. The positions in this plane were compared to those from synthetic colors given in the new Kurucz (1991) models to determine the value of $\log g$. These values were corrected using the equation in chapter 5 so that the gravities were on the scale of Edvardsson *et al.* (1993). Not all stars in a cluster had Stromgren photometry. In these cases, differential $\log g$ values (*i.e.*, values with respect to those cluster stars having Stromgren photometry) were determined using equation 16.3 of Gray (1992) and the $(B - V)$ data. For M 67, this approach is not so straight forward since the stars are near the turnoff, where $\log g$ values are double-valued for a given $(B - V)$. For this cluster, we derived differential $\log g$ values by comparison of the stars' positions in the color-magnitude diagram and the gravities of the Revised Yale Isochrones (Green, Demarque & King 1987). Logarithmic gravities are listed for each star in column three of Tables 13, 14, 15, and 17; column four of Table 16; column five of Table 18; and column six of Table 19.

For NGC 188, no Stromgren photometry was available for any of the stars in the sample presented here. Hence, we were forced to rely on the absolute gravity values from isochrones. Dr. P. Demarque kindly made available isochrones specially developed for comparison to the color-magnitude diagram of NGC 188. These isochrones were those discussed, utilized, and shown in Demarque, Green & Guenther (1992). I used the 7 Gyr isochrone with the LAOL opacities and the Grevesse solar mixture. This isochrone was shifted to roughly match the color-magnitude diagram of NGC 188 as constructed using photometry, kindly provided by Dr. K. Janes, of the high probability members according to Upgren *et al.* (1972). The vertical and horizontal shifts did not correspond to the reddening and distance modulus in Demarque *et al.* (1992) since the 7 Gyr was just slightly older than the best fit age they derived. The gravities estimated from the isochrones are given in column four of Table 20. On average, they are $\sim 0.15 - 0.20$ dex smaller than those estimated by Hobbs *et al.* (1990) based on an adopted distance modulus, stellar temperature, and stellar mass-to-luminosity ratio. However, the distance modulus adopted by them is ~ 0.6 mag smaller than that in Demarque *et al.* (1992). This means, all other things held constant, that Hobbs *et al.*'s $\log g$ values should be $\sim 0.20 - 0.25$ smaller than those of the isochrones. Therefore, once differences in distance modulus are accounted for, the values are reasonably consistent.

3.3 The Abundance Analysis

Abundances were derived using the LTE synthesis program RAI10 provided Dr. M. Spite. We used this software to construct curves of growth for the 7774 Å O I triplet. Abundances were read off from the curve of growth using the measured equivalent widths. Model atmospheres from the exten-

sive new grid of Kurucz (1991) were selected according to the star's/cluster's temperature, gravity, and metallicity. Microturbulent velocities, required by the abundance software, were calculated (as in chapter 5) using equation (9) of Edvardsson *et al.* (1993). We adopted the same O I triplet gf values as used in chapter 5 (*i.e.*, those of Wiese, Smith & Glennon 1966). The absolute abundances, $\log N(\text{O})$, were converted to the usual relative abundances $[\text{O}/\text{H}]$ by deriving the solar abundance from similar moderate-resolution spectra of the day-time sky. This was discussed in chapter 5 and we use the equivalent width of the UH 2.2-m faint object spectrograph (FOS) spectra to determine an abundance of 8.880 (on the usual scale where $\log N(\text{H}) = 12.0$). As noted in chapter 5, higher resolution spectra of the asteroid Vesta seem to indicate a slightly larger value 8.918, which is in excellent agreement with the meteoritic value of 8.93 (Grevesse & Anders 1989). Thus, it is *possible* (but cannot be definitively concluded) that, in the process of normalizing the absolute abundances to solar, systematic errors as high as ~ 0.04 dex have arisen. Fortunately, this potential effect is small.

In all the tables, estimated abundance uncertainties are presented in the columns next to the ones giving the O abundances themselves. These uncertainties are based on the individual uncertainties due to the estimated errors in the equivalent widths, T_{eff} , and $\log g$. The abundance uncertainties due to the errors in the equivalent widths were easily read off the curves of growth themselves. The equivalent width uncertainties were estimated from the individual spectra by making allowance for reasonable variations in the continuum setting and wavelength range of integration, rather than from more formal treatments based on S/N and resolution. Previous comparisons of our equivalent widths with others for halo stars and field Pop I stars (in chapters 2 and 5) indicate that these are likely to be overestimates of 1σ uncertainties. However, we treat them as such here in order to be conservative.

The many curves of growth we computed allowed us to quantify the uncertainty in abundance due to errors in T_{eff} and $\log g$. Typically (the uncertainties in abundances are not constants for an error of a given magnitude, but vary somewhat with other parameters), an error of ± 100 K in T_{eff} leads to an error of ∓ 0.08 dex in the O abundance. Similarly, an error of ± 0.5 dex in $\log g$ will lead to an error ± 0.15 dex in O abundance. Errors due to uncertainties in the adopted $[\text{Fe}/\text{H}]$ are negligible compared to the others, so were ignored. An uncertainty in $\log g$ of ± 0.2 dex was assumed for all stars. Temperature uncertainties were estimated from the dispersion in the values derived from the $(B - V)$, $(b - y)$, and β indices. If all of these were not available, an uncertainty of ~ 75 K was adopted unless we had reason to believe it was larger (as was the case for a few of the stars in the older clusters where there are some notable discrepancies in the photometry). All of the

uncertainties were added in quadrature to derive the final values given in Tables 13 – 20.

4. Results

4.1 Mean Cluster $[O/H]$ Abundances

Mean cluster abundances were calculated by computing the weighted averages (using an individual star's σ^{-2} value as the weighting factor) of the abundances given in Tables 13 – 20. This first involved eliminating some stars due to likely non-membership and stars that are decidedly too hot for the O I triplet to yield reliable abundances. First, we discuss membership. All Pleiades stars in Table 13 were retained as members and used in the weighted mean abundance determination. Based on its discrepant $[Fe/H]$ value from Laird *et al.* (1988) and the discrepant $[O/H]$ value we find here, Tr 23 was excluded as a Coma member in the determination of the cluster's mean abundance. The thorny issue of UMa group membership has recently been addressed by Soderblom & Mayor (1993). Based on the spectroscopic and kinematic information they provide, HR 3064 can be safely excluded as a group member, in agreement with their assessment. Similarly, it seems likely that HR 235, 7172, and 7451 can also be reasonably excluded as members. However, it is not so clear to us whether HD 151044 can be excluded as Soderblom & Mayor do. Therefore, we retain it as a member. The three remaining stars can be safely retained, we feel, as high likelihood members (again, in agreement with Soderblom & Mayor's assessment). Our $[O/H]$ abundances seem to confirm the low proper motion memberships for Praesepe stars JS 103 and JS 638; hence, they are excluded in the mean cluster abundance determination. All Hyades stars in Table 17 are retained as members. All NGC 752 stars in Table 18 are retained as members. S 762 is excluded from the mean $[O/H]$ determination for M 67 given its low proper motion and radial velocity membership probabilities; its $[O/H]$ abundance is somewhat, though not drastically, lower than the final mean cluster abundance. All NGC 188 stars in Table 20 were retained as members in the mean abundance determination.

It is well known that in hotter stars, significant NLTE effects become important for the O I 7774Å triplet. Chapter 5 demonstrated that such NLTE effects arise at effective temperatures greater than 6200 – 6300 K. Therefore, stars hotter than this must be excluded from the mean cluster abundance determination. In fact, this has already been done here by not including such stars in the Tables. Additional, hotter stars (roughly 5-10 in each of the clusters observed with Hydra, and 1 or 2 in the other clusters) were observed in several of the clusters in order to confirm the results of chapter 5. These results are confirmed and we do not detail the abundances here since they are not useful for our purposes. However, to give some sort of indication of the effect,

two of the hot stars (KW 227 and 250) in Praesepe were deliberately included as a demonstration (they were, of course, excluded in the mean abundance determination). Inspection shows that the $[O/H]$ abundances of these high probability cluster members are much larger than the other cooler members. Thus, given the results of chapter 5, we do not believe that further consideration of NLTE effects on the abundances are warranted for the stars whose abundances are presented here. Indeed, we do not see the expected effect, larger $[O/H]$ values for the hotter stars, if this were not the case. The final weighted mean $[O/H]$ abundances for each cluster are listed in column 4 of Table 21. Column five lists our estimate of the uncertainty in the abundance. This estimate is composed of both the error in the weighted mean abundance plus our assessment of possible systematic errors (due to the abundance normalization, errors in reddening, *etc.*).

4.2 Comparison with Other Results

4.2.1 Hyades & UMa

Garcia Lopez *et al.* (1993) have derived O abundances for Hyades and UMa group stars using a NLTE analysis. For $T_{\text{eff}} \leq 6450$ K, they find mean abundances of $[O/H] = -0.07$ for the Hyades and $[O/H] = -0.09$ for the UMa group. Comparison with our values in Table 21 reveals that there is a large discrepancy between our results and theirs. This seems to be due to the following: First, the solar abundance they derive from their NLTE computations is quite large at a value of $\log N(O) = 9.16$. This value is certainly not correct in an absolute sense as analysis of meteorites, LTE analysis of the solar 6300 and 6363 Å [O I] lines, and LTE analysis of the 7774 Å triplet all result in a consistent solar value of $\log N(O) \sim 8.9$. If, however, this were a systematic error in their analysis (*i.e.*, it affected the Hyades and UMa abundances in the same way), then the $[O/H]$ values (*i.e.*, incorrect object abundances normalized to an identically incorrect solar abundance) should still be accurate. Indeed, Garcia Lopez *et al.* suggest that their van der Waals broadening parameter may be too low— this would be such a systematic effect.

However, we are not so certain that this is the real cause of the discrepancy. The discrepancy may be due in part to their adopted solar O I equivalent widths, which are much higher than those noted in chapter 5; these latter values, all in good agreement, consisted of high-resolution observations of Vesta, moderate-resolution spectra of the daytime sky, and the measurements of Edvardsson *et al.* (1993). We also can not rule out that the discrepancy arises, in part, from the NLTE calculations. It is well known that NLTE calculations (*e.g.*, Kislman 1991 and Tomkin *et al.* 1992) are unable to reproduce the correct solar abundance. Therefore, since such NLTE analyses

are presently unable to yield the correct solar abundance, there is no reason to believe that they yield identically incorrect abundances for other stars of differing T_{eff} , $\log g$, and $[\text{Fe}/\text{H}]$.

Two lines of evidence may support this speculation: (1) Ignoring the Li dip region of the Hyades and UMa group, there is a clear systematic effect with T_{eff} in the Garcia Lopez *et al.* (1993) abundances in that larger $[\text{O}/\text{H}]$ values are derived with increasing T_{eff} —therefore we wonder if NLTE effects have truly been removed in their analysis. (2) If we take their Hyades equivalent widths for stars having $T_{\text{eff}} \leq 6300$ K and derive our own LTE abundances, we obtain larger *absolute* abundances than their NLTE abundances by ~ 0.14 dex on average. However, repeating the same exercise with the Sun, our LTE analysis yields an abundance ~ 0.12 dex *lower* than theirs. Therefore, this may be evidence that their NLTE analysis is not self-consistent; An admittedly small piece of possible confirming evidence that the Garcia Lopez *et al.* NLTE stellar abundances may be too small comes from the O abundances of Edvardsson *et al.* (1993), a subset of whose field stars Garcia Lopez *et al.* (1993) use to compare with their cluster results. For HR 1983, Edvardsson *et al.* derive an $[\text{O}/\text{H}]$ value of -0.14 , substantially larger than the value of -0.32 found by Garcia Lopez *et al.* (1993). The former’s abundances are on a scale of O abundance given by the 6300 Å [O I] line, which is thought to be free from NLTE effects.

We are also unable to reproduce the Garcia Lopez *et al.* (1993) LTE abundances using their equivalent widths and T_{eff} values. The LTE O abundances we found for the five stars of theirs we checked were ~ 0.2 dex smaller. This is even more surprising given that when our analysis was repeated using the value of the Van der Waals broadening, we found that their abundances should be ~ 0.05 dex *lower* than ours. We do not believe our abundances to be in error since the solar abundance we find from our software and model atmospheres is very close to the meteoritic abundance. Since Garcia Lopez *et al.* did not indicate a solar abundance determined using their LTE analysis procedures, we are unable to explain this discrepancy. We do note, that for both their Hyades and UMa sample, the LTE abundances do not increase with T_{eff} for $6300 \leq T_{\text{eff}} \leq 6700$. This is quite unexpected given the results in chapter 5. We also found that we were unable to reproduce the O abundances of Clegg, Lambert & Tomkin (1981) for their stars having $T_{\text{eff}} \geq 6200$ K. Again, our abundances for the hotter stars are lower than theirs. However, even though they do not derive an LTE solar abundance either, we can state in this case that abundance differences are *not* systematic. For stars having $T_{\text{eff}} \leq 6100$ K, we are able to reproduce the abundances they derive.

In sum, we are skeptical of the Garcia Lopez *et al.* (1993) abundances. The bottom line is that LTE models reproduce the solar abundance very well (with no large enhancement of the usual fudge factors such as C_6). Chapter 5 directly showed that LTE analysis yields accurate O abundances from

the 7774 Å triplet when applied to dwarfs having $T_{\text{eff}} \lesssim 6300$ K. Since our Hyades and UMa stars meet this requirement, we believe that our abundances are accurate. We also note one additional item of interest here. The weighted mean abundance of the Garcia Lopez *et al.* stars *reanalyzed* here is slightly larger than the weighted mean abundance determined from the remaining stars in Table 17 ([O/H] of +0.31 *vs.* 0.25). First, we alert the reader that an intermediate value (+0.28) was adopted as the final abundance. While the difference is not formally significant given the errors, we speculate whether it could be due to differences in the equivalent width scale. While our observations of VB 66 yield essentially the same equivalent width, their value for HR 2047 is substantially larger than ours. While this alone is not any rigorous indication of systematic differences in equivalent width scales, we feel this issue must be mentioned since all the Hyades observations of theirs that we have reanalyzed were obtained with what appears to be very similar or identical instrumentation (same telescope, spectrograph, detector) as the O I data of Abia & Rebolo (1989). These latter data have been shown to have equivalent widths which are systematically too large (see chapter 2 and Tomkin *et al.* 1992).

4.2.2 M 67

Cohen (1980) derived O abundances for four M 67 giants using the 6363 Å [O I] line. The mean abundance, corrected for amount of O locked up in CO, is [O/H] = -0.09 (though it is not clear to us that this value is consistent with the four values, given in her Table 6B, which are all positive or zero). With her assumption of [Fe/H] $\sim$ -0.4, her value for [O/Fe] is $\sim$ 0.3 dex. We tend to place a bit more faith in the O abundances of Brown (1985), who derived the O abundance from the 6300 Å line in three M 67 giants. As he notes, the 6300 Å equivalent widths are larger than those of the 6363 Å line; furthermore, comparison with CCD data (Cohen 1983) has revealed that the equivalent widths of Cohen (1980) from photographic plates may suffer from systematic effects. Brown's (1985) value is [O/H] $\sim$ -0.2. We note that, if the [Fe/H] value of Hobbs & Thorburn (1991) is to be preferred over Brown's assumed value for M 67, then the [O/H] ratio may need to be raised slightly ($\sim$ 0.05 - 0.08 dex). If the value of [O/H] $\sim$ -0.2 is correct, then M 67 would fall exactly in place with the other older clusters in the [O/H]-Age plane (discussed below). However, we must ask if Brown's value truly represents the initial (or main-sequence star) O abundance. We will address this important question later.

Dr. A. Boesgaard has kindly communicated some preliminary results from work by her and Dr. E. Friel on O abundances of M 67 dwarfs derived from high resolution spectroscopy. She informs

the author that their data indicate that M 67 is slightly O enhanced relative to its $[\text{Fe}/\text{H}]$ value. Given $[\text{Fe}/\text{H}] = +0.02$ from Friel & Boesgaard (1992), we interpret this as excellent agreement between their unpublished O abundances and those found here.

4.3 Cluster Ages

The logarithm of the cluster ages (in years) are listed in column 8 of Table 21. For the younger clusters, ages were taken from Boesgaard (1989) or Boesgaard & Friel (1990) essentially unchanged. For NGC 752, we adopted the canonical age of 2 Gyr. For M 67 the two age estimates of Demarque *et al.* (1992) and Hobbs & Thorburn (1991) are available. While we acknowledge that the spectroscopic T_{eff} values of the latter help avoid some uncertainties in the comparison with stellar evolutionary models, we have given the Demarque *et al.* (1992) higher weight and adopt an age of 4.2 Gyr. Two things to mention concerning this are that: 1) we do not believe that actual precision to 0.1 Gyr can be achieved by current dating methods and 2) we have neglected the 0.01 mag difference between the reddening used by Demarque *et al.* (1992) and us (given the uncertainties, this is of little consequence, but it must be acknowledged that the adopted age is not strictly self-consistent with our reddening). Similarly, we have the age estimates for NGC 188 of Demarque *et al.* (1992) and Hobbs *et al.* (1990). Giving the former value slightly higher weight, we adopt an age of 7.0 Gyr. Realistic, but admittedly subjective, estimates of the uncertainties in the ages are given in column 9 of Table 21. These do not encompass all possible values of the cluster ages, but are rather meant to represent, in our assessment, conservative 1σ uncertainties.

We emphasize at this point that the ages we adopted are those commonly found from evolutionary models that do not assume convective overshoot. In recent years, much evidence has been put forth to argue that conventional treatments of convection are inadequate. Models which include convective overshoot (*e.g.*, Meynet, Mermilliod & Maeder 1993) lead to younger ages (by factors of $\sim 30 - 100\%$) for intermediate age clusters. It is not the intent of this paper to review the evidence for and against convective overshoot. Nor are we going to enter the debate about which assumption is correct. Rather, we think it at least remains to be demonstrated that ages from standard models are incorrect when new opacity data is utilized in the models (see Stothers & Chin 1992) and we merely inform the reader that we are adopting ages in accord with standard models.

5. Discussion

5.1 Age-[Fe/H] Relations

Figure 31 shows the Fe abundance of our clusters plotted against their ages. As has been noted before (Boesgaard 1989, Nissen 1988, Friel & Janes 1993, Geisler *et al.* 1992), there is clearly no obvious global age-[Fe/H] relation for this sample (or the larger samples of the latter two references) of clusters. This result differs with the age-[Fe/H] relation for solar neighborhood field stars found by Twarog (1980). Thus, one reaches the usual conclusions that either: 1) the clusters and field stars have shared the same chemical evolution history, but the relation found for one is incorrect or 2) the clusters and field stars have not shared a common chemical evolution history— *i.e.*, there have been temporal and/or spatial variations in the IMF and SFR or the disk “mixing” timescales are different for the clusters and field stars. The first option has been argued by Carlberg *et al.* (1985) and Barry (1988), whose results show no real age-[Fe/H] relation for the field stars. However, the stronger temporal Fe abundance gradient of Twarog (1980) seems to be consistent with the recent results of Rana & Basu (1992) and Twarog (1980) discusses the problems with using clusters to sample the history of Galactic chemical evolution history.

In their recent tome on chemical evolution in the Galactic disk, Edvardsson *et al.* (1993) state that “There is a considerable variation in metallicities of stars formed at a given time in the disk, which means that there is only a weak correlation between age and metallicity”. However, if one were to bin the data in their figure 14a, the resulting relation would not be unlike that found by Twarog (1980). Therefore, the issue of age – [Fe/H] relations is, in large measure, one of semantics as well as statistics. While there is no exact one-to-one relation (*i.e.*, the fact that there exist some stars with higher Fe abundances than stars that are younger and there exist some stars with lower Fe abundances than stars that are older), globally there may be such a trend (*i.e.*, upon binning the data to search for mean trends). We will not directly enter into this debate here, but instead will see what light can be shed on the matter using our O abundances.

5.2 Age-[O/H] Relations

Figure 32 shows our clusters’ O abundances plotted against the ages from Table 21. This plot exhibits a clear global trend between age and [O/H] ratio. Indeed, except for the Hyades and M 67, there is a clear monotonic relation between the two variables. If one adopts the canonical relation $[O/H] = +0.5 \times [Fe/H]$ for solar neighborhood field stars and applies this to Twarog’s (1980) sample, the resulting [O/H] – age relation is similar to the clusters; however, a slope of +0.7 rather than +0.5 gives a better match. Additionally, the global trend of the data in $[\alpha/H]$ – age plot in figure 14b of Edvardsson *et al.* (1993) is in fair agreement with the cluster data. However, two things must

be noted about this latter correspondence: 1) the α elements in this plot include only Ca and Si 2) even if their O abundances did follow this same trend, we do *not* take this as proof of a similar chemical evolution between the clusters and field stars given possible discrepancies between the O abundances of Edvardsson *et al.* (1993) and our abundances as discussed in chapter 5.

The lack of any relation between $[\text{Fe}/\text{H}]$ and age might be a bit surprising if the timescale for production of Fe by Type Ia supernovae is $\leq 10^9$ yr (see Smecker-Hane & Wyse 1992). On the other hand, the relation between age and O abundance may not be particularly surprising if the timescale for production of O by Type II supernovae is $\sim 10^8$ yr (see Gilmore, Wyse & Kuijken 1989; Smecker-Hane & Wyse 1992). If, as suggested by Smecker-Hane & Wyse (1992), most Type Ia systems are CO-He white dwarfs, which have Fe production timescales of $10^9 - 10^{10}$ yr, then the behavior of Fe abundance with age is understandable. Additionally, the large scatter demonstrated by field stars in the $[\text{Fe}/\text{H}]$ vs. age plane (*e.g.*, Edvardsson *et al.* 1993), is also consistent with a long timescale for Fe production. Since the Type Ia Fe production timescale constrains the duration of halo formation, the longer timescales of the CO-He systems suggest a slow (few Gyr) halo collapse which would accommodate alleged age spreads among Galactic globular clusters and field stars (Gilmore *et al.* 1989, Smecker-Hane & Wyse 1992).

5.3 $[\text{O}/\text{H}]$ versus $[\text{Fe}/\text{H}]$

Figure 33 shows the cluster $[\text{O}/\text{H}]$ values plotted against the $[\text{Fe}/\text{H}]$ values. These can be compared to the trend shown by field stars using figure 4 of Garcia Lopez *et al.* (1993). Given the uncertainties, the cluster values cannot be said to be inconsistent with the behavior (decreasing $[\text{O}/\text{H}]$ with decreasing $[\text{Fe}/\text{H}]$) shown by the field stars. However, the clusters studied here do not, by themselves, exhibit a significant trend. The statistical significance is at the $\sim 70\%$ level, though in the same sense as that seen in field stars.

5.4 A New Perspective

Figure 34 shows the cluster $[\text{O}/\text{Fe}]$ values plotted against their ages. The mean trend seen is for larger $[\text{O}/\text{Fe}]$ ratios at younger ages. This is the opposite of what is seen in field stars having an assumed age - $[\text{Fe}/\text{H}]$ relation like Twarog's (1980). It is well known that field stars show increasing $[\text{O}/\text{Fe}]$ ratios with decreasing $[\text{Fe}/\text{H}]$ (hence, increasing age with the assumption of an age - $[\text{Fe}/\text{H}]$ relation). Certainly this difference is not surprising given the very different behavior between age and $[\text{Fe}/\text{H}]$ for the clusters and solar neighborhood field stars.

The difference may be surprising, however, in light of the results of Friel & Janes (1993), who find that their Mg line strengths are stronger relative to Fe for older clusters than for younger clusters. If interpreted as an abundance effect, this is similar to the behavior inferred for field stars if one assumes that an Fe – age relation exists for them (*i.e.*, $[\text{Mg}/\text{Fe}]$ is seen to increase with decreasing $[\text{Fe}/\text{H}]$ in field stars). Since the chemical evolution of Mg and O are expected to be quite similar, we might hope to find the same behavior in our cluster data. Figure 35, though, indicates that the O abundances are highest relative to iron for the younger clusters. Friel & Janes (1993) note that the effect they see may also be due to surface gravity. However, this solution is unattractive in light of the present results since Friel & Janes indicate that surface gravity effects fail by a factor of 2 – 3 to fully account for the trend they see, let alone overcome an effect in the opposite direction as seen here. Clearly, there remains much work to be done concerning the chemical evolution history of elements other than Fe in the Galactic disk. In any case, our results shown in Figure 36 (described next) are consistent with those of Friel & Janes (1993), who find that the $[\text{Mg}/\text{Fe}]$ anomaly is correlated with age only and not with $[\text{Fe}/\text{H}]$.

The relation between the clusters' $[\text{O}/\text{Fe}]$ and $[\text{Fe}/\text{H}]$ ratios are shown in Figure 36. This is little more than a scatter diagram, the data having a correlation coefficient of only $r = 0.006$, and is not very useful in gathering information in the same sense done with field stars (where there is a clear increase in $[\text{O}/\text{Fe}]$ with decreasing $[\text{Fe}/\text{H}]$ for $[\text{Fe}/\text{H}] \geq -1.0$) where $[\text{Fe}/\text{H}]$ is treated as a chronometer. As noted in the introduction, Wheeler *et al.* (1989) suggest that $[\text{O}/\text{H}]$ may serve as a more useful chronometer in terms of studying Galactic chemical evolution. They further suggest that the corresponding diagram of $[\text{Fe}/\text{O}]$ *versus* $[\text{O}/\text{H}]$ may give a different perspective to the situation while containing no new information *per se*. This suggestion is seemingly verified in Figure 37, which shows the cluster $[\text{Fe}/\text{O}]$ ratios plotted against $[\text{O}/\text{H}]$. Couching the data, which have a correlation coefficient of $r = -0.91$, in this manner strongly suggests that $[\text{O}/\text{H}]$ is a more useful chronometer with which to study the chemical evolution and history of open clusters.

5.5 The Present Day O Abundance

5.5.1 The Present Day O Abundance Problem

A well-known problem concerning the Galactic chemical evolution of oxygen is the fact that the O abundance of the Sun and metal-rich, nearby field stars is much larger (by ≥ 0.2 dex) than abundances identified as representative of the present day ISM. O abundance values derived from H II regions are typically 0.3 dex lower than the solar/stellar values (Osterbrock, Tran & Veilleux

1992). These lower abundances are seemingly confirmed by observations of gaseous nebulae (Peimbert, Storey & Torres-Peimbert 1993), planetary nebulae (de Freitas Pacheco 1993), and (young) supergiants (Luck & Lambert 1985). These results seem to indicate that there has *not* been an increase in O abundance in the disk during the past 4.6 Gyr. The two major mechanisms proposed to explain this behavior are: a) infall of halo gas having abundances characteristic of those produced by Type I supernovae and/or b) an increase, in the disk, of Type I supernovae compared to Type II supernovae during the past 5 Gyr or so (de Freitas Pacheco 1993).

5.5.2 *The Open Cluster Results*

The O abundances in our younger clusters clearly indicate that there has been an increase in [O/H] over the past 4.6 Gyr. The results for the Pleiades, the Hyades, Coma, and Praesepe would seem to indicate that the present day ISM O abundance is ~ 0.2 dex higher than the solar abundance. Thus, the cluster abundances further aggravate the controversy surrounding the present day O abundance. We now try to resolve this discrepancy and see what the implications of such a resolution are.

5.5.3 *Resolving the Present Day O Abundance Discrepancy*

First, we consider the H II region O abundances. Of particular concern here is the presence of possible temperature fluctuations along the line of sight (within the H II regions) which could lead to incorrect abundances. Such temperature variations can apparently explain the O abundance differences between the H II regions studies and the solar/stellar studies (Walter, Dufour & Hester 1992). Thus the reliability of the H II region O abundances may be less than stellar photospheric abundances.

de Freitas Pacheco (1993) has recently derived O abundances for a sample of 122 planetary nebulae. It is found that the average O abundance is some 0.2 dex lower than the solar value. We emphasize, however, that an implicit and explicit assumption in this analysis is that the planetary nebulae progenitors are *not* affected by mixing and burning of O— *i.e.*, that their O abundances are precisely equal to the abundance of the ISM out of which their progenitors formed. Is there any evidence to indicate the contrary? We believe there are two pieces of direct evidence which indicate otherwise. First are the abundances of M 67 giants derived by Brown (1985), which are some 0.2 dex lower than the O abundances found here from dwarfs (which, recall, seem to be further confirmed by the unpublished values of Boesgaard & Friel). These results could indicate that there has been O $\rightarrow$ N processing of the primordial material in these giants.

The other piece of direct evidence that there has been O→N cycling comes from the recent study of CNO abundances in three Population I A supergiants by Venn (1993). She finds [O/H] abundances in her three objects of -0.19 , $+0.02$, and -0.04 dex. These seemingly confirm the constant or decreasing O abundance in the Galactic disk over the past 4.6 Gyr. However, she finds startling overabundances of N in her stars, which have [N/O] of $+0.98$, $+0.93$, and 0.76 dex. These may be compared to the value of $+0.50$ predicted by the calculations of Maeder & Meynet (1988). If we enforce number conservation and assume that all additional N comes from O→N cycling, then the corrected [O/H] abundances are in very good agreement with those seen in our youngest clusters. Hence, it may well be that O abundances determined from either planetary nebulae or supergiants do *not* reflect the initial O abundance that their progenitors formed with.

A more difficult obstacle to making the determination that the stellar O abundances are to be preferred over other determinations is the recent work by Peimbert *et al.* (1993). They present O abundances for NGC 6572, M 17, and Orion which are based on recombination lines that are independent of possible line-of-sight temperature variations noted above. The final abundances are still subsolar. One possibility may be that these observations are of *gaseous* nebulae whereas most of the ISM O may be locked up in grains (Meusinger 1992). The usual objection to this possibility is the constancy of Ne-to-O from nebula to nebula. However, whether this constancy is preserved in light of the possibility of line-of-sight temperature variations is not clear. Clearly, this specific issue concerning oxygen and the dust-phase ISM merits further study.

5.5.4 Implications of the Young Open Cluster O Abundances

We now ask what are the implications of the young cluster O abundances, if taken at face value. The first implication of the young cluster O abundances would be that the stellar model calculations of Maeder & Meynet (1988) and Renzini & Voli (1981) may be incorrect. As noted by Venn (1993), the Maeder & Meynet (1988) evolutionary and dredge-up calculations do not agree with her observed [N/O] ratios. de Freitas Pacheco (1993) states that the maximum depletion of surface O abundance in likely planetary nebulae progenitors is ~ 0.1 dex. As noted above, this is not consistent with the cluster O abundances.

Venn (1993) notes that one possibility is that additional mixing, beyond that included in the models, takes place in the supergiants (or the planetary nebulae progenitors). Particularly interesting in this regard are the Yale rotational models (*e.g.*, Sofia, Pinsonneault & Deliyannis 1991). These models have met with good success in the comparison with observations, including: prediction of

the Li “dip” in F stars, accounting for observed spreads of Li abundance in open clusters, accounting for the apparent dispersion in Li abundances in halo stars, agreement with the observed rotation velocities of halo horizontal branch stars compared with those of halo dwarfs, and the ability to explain Be-deficient stars that still retain measurable Li. We speculate, as have others, that this additional mixing (not included in standard modeling) may be able to explain the observed dispersion in globular cluster giants’ O abundances (Kraft *et al.* 1992, Sneden *et al.* 1991), though it must be acknowledged that it is not clear yet how much of a role primordial variations play in producing such a spread. Indeed, it is interesting that the spread of planetary nebulae O abundances seen in figure 1 of de Freitas Pacheco (1993) is quite similar to that observed in globular cluster red giants.

If this additional mixing is occurring, then there are two major implications: 1) The most relevant number, in terms of the present day ISM O abundance, is the maximum abundance seen in figure 1 of de Freitas Pacheco (1993). This value, $[O/H] \sim 9.1$, is essentially the same (to within the uncertainties) as that derived here from the cluster stars. 2) Halo star O abundances derived from red giants may be affected by additional mixing. Therefore, the most relevant number in terms of the halo $[O/Fe]$ ratio may be the upper envelope in figure 11 of Kraft *et al.* (1992). This value, $[O/Fe] \sim 0.5$, is in very good agreement with the mean value (0.53) found in chapter 2 for halo dwarfs, whose O abundances were determined from the permitted 7774 Å O I triplet. We strongly urge further investigation of the Yale models and their implications for O abundance processing in both metal-rich and metal-poor stars having a range of masses. A final (and obvious implication) of the open cluster O abundances is that since planetary nebulae and H II regions are used to derive O abundances of extragalactic systems and to deduce the O abundance gradient in the Galactic disk, if their observed O abundances do not represent their initial O abundances or account for O locked up in dust grains, then inferences made about such gradients or extragalactic abundances may be incorrect.

5.6 Closing Thoughts and Future Work

Given the small cluster sample size, further analysis and interpretation at this point is not particularly worthwhile. Issues such as “metallicity” gradients (using O as a metallicity indicator) with height above the plane or with Galactic radius will require a larger cluster sample with an extended baseline of these parameters. Independent confirmation of the results found here is needed. Additionally, several other clusters could be observed to extend and fill in the age baseline. Good

candidates include α Per¹³, which is younger than the Pleiades; M 11, which is intermediate in age between the Pleiades and the Hyades; Mel 66, which is a candidate for being older than NGC 188; and NGC 3680, a cluster with an age probably between M 67 and NGC 752. Observations of interesting, more distant clusters such as Mel 66 or NGC 2420, will be quite difficult using the 7774 Å O I triplet due to the forest of OH emission lines in the vicinity. With the next generation of large (8-m, 10-m, and 25-m) telescopes it should be possible to secure observations of the 6300 Å O I line (assuming that accommodation is made for the required high resolution, efficient spectrographs). It should be stressed that if the 6300 Å line is to be used, then the O abundance analysis should be accompanied by a complementary one to determine the “metallicity” of the cluster— where, in this case, “metallicity” refers to the abundance of those elements which are the predominant electron donors (Fe, Si, and the like) in the stellar atmospheres. Such a determination is necessary given the sensitivity of the [O I] abundances to such overall metallicity. Spectroscopic determination of parameters is also urged to reduce the attendant uncertainties of reddening values and reliance on stellar models.

Of particular interest to observe would be the very metal-poor ($[\text{Fe}/\text{H}] \sim -1$), as measured by Geisler *et al.* 1992), clusters. Particular questions to be addressed besides those noted above are: 1) does the [O/H] ratio maintain an increase at young age or does it scatter like the [Fe/H] ratio? 2) Are there more clusters such as the Hyades and M 67 which seem to lie off from the global relation? 3) What are the [O/H] ratios in the $[\text{Fe}/\text{H}] \sim -1$ clusters and are they more similar to the [O/H] ratio (-0.5) of field stars of the same [Fe/H] or more similar to other clusters with similar ages but larger values of [Fe/H]? Once more clusters have O abundance determinations, questions such as these and the issues addressed by Friel & Janes (1993) and Geisler *et al.* (1992) can be more closely examined.

In the meantime, only a few broad conclusions can be reached: 1) The open clusters studied here show a clear global relation between O abundance ([O/H]) and age. 2) This behavior is very different than what is seen using Fe abundance, which shows no relation to the age of the cluster. 3) This difference may be due to CO-He white dwarfs comprising the majority of Type Ia supernovae. 4) The [O/H] ratio and the [Fe/O] *versus* [O/H] plane offers a very different (and more useful) perspective, as suggested by Wheeler *et al.* (1989), of the history of open cluster chemical evolution. 5) Hence, O abundance may provide an improved perspective in viewing the chemical evolutionary history of the Galactic disk’s *field* star population. A spectroscopic study of a carefully selected

¹³ One has to be careful about observing the O I triplet in quite young stars that exhibit classical T Tau effects. In these cases, the triplet may no longer be a useful abundance indicator.

sample of stars— such as that presented by Twarog (1980)— might prove invaluable in studying the chemical evolution of the field population and settling the question as to the existence and nature of an age — “metallicity” relation. Additionally, such an analysis might also include spectroscopic Fe abundance determinations in an effort to repeat Twarog’s (1980) analysis; a spectroscopic effort could further ensure that chemically peculiar stars (such as Am or Fm stars) get weeded out from the final sample. In any case, the initial open cluster results of this study seem to indicate that arguments rejecting O as a usable chronometer (arguments such as its sensitivity to IMF, SFR, *etc.* and arguments founded on retaining the Fe abundance as the traditional chronometer to maintain the status quo) may not stand up to observational evidence. 6) Our cluster O abundances imply that the present day ISM O abundance is supersolar and much larger than abundances measured in H II regions, planetary nebulae, and early-type stars. Given the O abundances derived for M 67 giants and the recent [N/O] overabundances found by Venn (1993), we suggest that stellar models which do not include additional mixing (such as that prescribed in the Yale rotational models) may be incomplete. If this is the case, then it is argued that a different interpretation of the existing planetary nebulae and supergiant O abundances is consistent with the abundances determined from young open clusters. Furthermore, use of average [O/Fe] ratios in field halo giants may underestimate the true initial [O/Fe] ratios. In any case, the O abundance of disk population stars richly deserves to be investigated further and, if warranted by the findings, to complement (if not supersede) the Fe abundance in the study of the history of chemical evolution in the Galactic disk.

Table 13. Pleiades Parameters and Abundances

Star	T_{eff} K	$\log g$	ζ km/s	EW mÅ	σ mÅ	[O/H]	σ dex
HII 739	5901	4.35	1.35	264	12	+0.31	0.10
HII 761	5792	4.37	1.25	210	19	+0.32	0.15
HII 923	5954	4.34	1.40	245	18	+0.17	0.13
HII 948	5964	4.17	1.65	283	14	+0.25	0.10
HII 1101	6008	4.33	1.50	299	10	+0.36	0.09
HII 2506	6083	4.34	1.50	281	25	+0.22	0.16

Table 14. Coma Parameters and Abundances

Star	T_{eff} K	$\log g$	ζ km/s	EW mÅ	σ mÅ	[O/H]	σ dex
Tr 23 ^a	5967	4.33	1.45	153	22	-0.36	0.16
Tr 85	5951	4.34	1.40	250	12	+0.19	0.11
Tr 102	5914	4.37	1.35	232	19	+0.15	0.17
Tr 132	5641	4.36	1.15	193	20	+0.23	0.19

^aCluster non-member ?

Table 15. UMa Moving Group Parameters and Abundances

Star	T_{eff} K	$\log g$	ζ km/s	EW mÅ	σ mÅ	[O/H]	σ dex
HR 235	6195	4.29	1.70	340 ^b	20	+0.34	0.11
HR 2047	5900	4.36	1.35	256	20	+0.26	0.14
				280 ^b	10	+0.39	0.08
HR 3064	5942	4.29	1.50	218 ^b	10	+0.01	0.09
HR 3391	5827	4.39	1.25	238 ^b	20	+0.26	0.15
HD 115043	5820	4.30	1.35	253	22	+0.31	0.15
HD 151044	6163	4.27	1.70	299 ^b	15	+0.19	0.10
HR 7172	6081	4.01	1.95	374 ^c	5	+0.41	0.07
HR 7451	6305	4.23	1.85	380 ^b	15	+0.39	0.08

^bO EW from Garcia Lopez *et al.* (1993).

^c7774 Å EW from Coude spectra described in Chapter 5.

Table 16. Praesepe Parameters and Abundances

Star	Other	P %	T_{eff} K	$\log g$	ξ km/s	EW mÅ	σ mÅ	[O/H]	σ dex
KW 181	JC 160	99	5992	4.28	1.55	211	12	-0.08	0.10
JS 103 ^d	Ak 388	80	6516	4.20	2.05	242	14	-0.39	0.10
KW 227	JS 324	99	6607	4.27	2.05	508	15	+0.64	0.09
KW 250	JS 334	99	6411	4.18	2.00	498	30	+0.67	0.19
KW 335	JS 388	99	5785	4.35	1.25	194	11	+0.06	0.09
KW 399	JS 435	98	5868	4.33	1.35	193	10	-0.02	0.08
KW 432	JS 464	99	5791	4.34	1.30	210	14	+0.15	0.11
JS 638 ^d	AK 1646	83	5847	4.33	1.35	273	21	+0.45	0.14

^dCluster non-member ?

Table 17. Hyades Parameters and Abundances

Star	T_{eff} K	$\log g$	ζ km/s	EW mÅ	σ mÅ	[O/H]	σ dex
VB 15	5772	4.31	1.30	254	14	+0.38	0.12
VB 31	6094	4.26	1.65	326 ^e	10	+0.37	0.09
VB 48	6207	4.28	1.70	367 ^e	20	+0.44	0.12
VB 61	6245	4.28	1.75	360 ^e	20	+0.38	0.12
VB 65	6118	4.24	1.70	344 ^e	20	+0.41	0.14
VB 66	6103	4.28	1.60	285	12	+0.19	0.08
VB 73	5914	4.22	1.55	271	13	+0.27	0.10
VB 88	6174	4.27	1.70	336 ^e	20	+0.34	0.12
VB 97	5887	4.28	1.45	250	12	+0.22	0.10
VB 105	6081	4.30	1.55	283 ^e	20	+0.21	0.14

^eO EW from Garcia Lopez *et al.* (1993).

Table 18. NGC 752 Parameters and Abundances

Star	T_{eff} K	$\log g$	ζ km/s	EW mÅ	σ mÅ	[O/H]	σ dex		
H 80	99	97	6061	4.35	1.50	234	18	+0.01	0.15
H 87	99	97	6134	4.33	1.60	215	20	-0.16	0.15
H 120	97	96	6061	4.35	1.50	190	22	-0.21	0.17
H 146	98	91	5770	4.39	1.20	138	20	-0.26	0.18
H 161	84	96	6025	4.36	1.45	196	25	-0.15	0.17
H 183	99	94	5666	4.43	1.10	124	22	-0.23	0.21
H 185	98	98	6243	4.31	1.70	268	25	-0.01	0.16
H 237	99	88	5998	4.36	1.45	204	17	-0.10	0.15
H 252	99	95	6098	4.34	1.55	200	23	-0.18	0.17

Table 19. M 67 Parameters and Abundances

Star	P_S %	P_μ %	$P_{\mu r}$ %	T_{eff} K	$\log g$	ζ km/s	EW mÅ	σ mÅ	[O/H]	σ dex
S 750	70	96	97	5846	4.12	1.60	92 ^J	12	+0.16	0.19
S 757	95			6013	4.21	1.65	86 ^J	8	−0.04	0.16
S 758	96			6077	4.30	1.55	256	29	+0.09	0.18
S 760	96			6168	4.21	1.75	232	22	−0.16	0.15
S 762 ^a	0	0	1	6153	4.20	1.75	237	31	−0.13	0.18
S 967	94	99	99	6189	4.20	1.80	283	23	+0.06	0.15
S 976	95			6000	3.99	1.90	254	34	+0.01	0.20
S 1007	94	99	100	6095	4.16	1.75	270	32	+0.07	0.18
S 1051	94	0	0	6080	4.32	1.55	261	25	+0.11	0.16
S 1062	96			6181	4.25	1.70	250	31	−0.06	0.19
S 1076	94			6226	4.21	1.80	231	31	−0.21	0.19
S 1234 ^g	95			6192	4.35	1.60	280	20	+0.12	0.14
S 1240	93	98	99	6180	4.29	1.65	276	27	+0.08	0.17
S 1256	94	98	99	5873	4.20	1.55	226	29	+0.09	0.19

^aCluster non-member ?

^JEW for 7771.9 Å line only.

^gNoted as SB in Mathieu *et al.* (1986).

Table 20. NGC 188 Parameters and Abundances

Star	P_{UMK} %	T_{eff} K	$\log g$	ζ km/s	EW mÅ	σ mÅ	[O/H]	σ dex
I-37	93	5800	3.89	1.90	180	23	-0.24	0.19
I-45	96	5885	3.92	1.90	186	20	-0.27	0.17
I-48	92	5727	3.88	1.85	178	19	-0.18	0.15
I-49		5940	4.32	1.45	156	25	-0.31	0.20
I-70	94	5713	3.87	1.85	156	18	-0.30	0.15
I-71	95	5936	3.94	1.95	191	28	-0.27	0.21
I-76	95	5973	3.95	1.95	218	23	-0.17	0.16
I-93	97	5936	3.93	1.95	215	29	-0.16	0.18
I-99	96	5859	3.91	1.90	168	19	-0.35	0.15
I-101	88	6042	3.97	1.95	185	18	-0.38	0.14
I-108	94	5896	3.93	1.90	201	27	-0.19	0.18
II-52	96	5676	3.86	1.80	127	18	-0.42	0.17
II-60	90	5829	3.90	1.90	179	21	-0.27	0.16
II-64	95	5881	3.92	1.90	200	21	-0.20	0.16
II-138	94	6299	4.0	2.15	229	24	-0.37	0.16
II-197	95	5867	3.91	1.90	198	22	-0.19	0.16
III-44	96	6123	4.05	1.95	228	21	-0.20	0.15
III-106	95	6046	3.97	2.00	198	18	-0.33	0.14

Table 21. Cluster Abundances and Parameters

Cluster	[Fe/H]	σ dex	[O/H]	σ dex	[O/Fe]	σ dex	log Age (yr) log yr	σ log yr
Pleiades	-0.04	0.06	+0.29	0.08	+0.33	0.10	7.85	0.15
Coma	-0.07	0.05	+0.18	0.11	+0.25	0.12	8.70	0.20
UMa	-0.09	0.07	+0.24	0.10	+0.33	0.12	8.48	0.20
Praesepe	+0.08	0.05	+0.02	0.08	-0.06	0.09	8.90	0.15
Hyades	+0.12	0.05	+0.28	0.10	+0.16	0.11	8.83	0.15
NGC 752	-0.09	0.07	-0.13	0.12	-0.04	0.14	9.30	0.20
M 67	-0.04	0.12	+0.03	0.13	+0.07	0.18	9.63	0.15
NGC 188	-0.12	0.16	-0.27	0.13	-0.15	0.21	9.85	0.10

Figure 31. Samples of the open cluster spectra presented and analyzed in chapter 6. Spectra of H II 948 (Pleiades) taken with the FOS on the UH 2.2-m, HR 7172 (UMa group) taken with the Coude spectrograph at the UH 2.2-m, and H 261 (NGC 752) taken with Hydra on the KPNO 4-m are shown in the top, middle, and bottom panels. H II 948 has one of the lower S/N values of the FOS spectra. The high resolution of HR 7172 spectra is not more obvious because the line profiles are rotationally broadened. H 261 is an example of one of the hotter (6480 K) stars whose O abundances were not derived in chapter 7 due to large NLTE effects.

Figure 32. Iron abundance ($[\text{Fe}/\text{H}]$) is plotted against the clusters' logarithmic ages for the open clusters in our sample.

Figure 33. Oxygen abundance ($[O/H]$) is plotted against the clusters' logarithmic ages for the open clusters in our sample.

Figure 34. Oxygen abundance ($[O/H]$) is plotted against iron abundance ($[Fe/H]$) for the open clusters in our sample.

Figure 35. [O/Fe] ratio is plotted against the clusters' logarithmic ages for the open clusters in our sample.

Figure 36. $[O/Fe]$ ratio is plotted against the iron abundance ($[Fe/H]$) for the open clusters in our sample.

Figure 37. [Fe/O] ratio is plotted against oxygen abundance ([O/H]) for the open clusters in our sample.

Chapter 7. Conclusions, Summary Comments, and Future Work

1. Main Conclusions

We now summarize the conclusions reached here and discuss these findings in a broader sense than done in each individual chapter. A very fundamental conclusion, reached in chapter 5, is that the 7774 Å O I triplet seems to yield abundances in agreement with those determined from the 6300 Å [O I] line for relatively metal-rich dwarfs having $T_{\text{eff}} \leq 6200\text{--}6300$ K. This is contrary to the findings of Nissen & Edvardsson (1992) and contrary to the claims made by others that the 7774 Å triplet may not yield reliable abundances due to NLTE departures or the effects of atmospheric thermal inhomogeneities. We strongly emphasize once again that it is *not* denied here that the 7774 Å triplet is formed in NLTE or that it is subject to atmospheric inhomogeneities. Indeed, it is recognized that analyses of the O I triplet in the Sun clearly indicate that the lines are formed in NLTE and that atmospheric inhomogeneities are manifest in the Solar line profiles themselves. The relevant question, which should not in any way diminish the analyses of, *e.g.*, Altrock (1968) or Kiselman (1991), in the author's mind is whether these NLTE departures or atmospheric inhomogeneities have an *observable effect* on O abundances derived from *flux* spectra. In the case of the Sun, the answer seems to be no. The Solar O abundance derived from LTE analyses of Solar flux spectra of the 7774 Å triplet is in very good agreement with the meteoritic value. This agreement exists despite the notable disagreement between LTE models and Solar intensity spectra at small μ and despite the line profile asymmetries, which are probably caused by thermal inhomogeneities.

We also noted that there seem to exist significant uncertainties in O abundances derived from the 6300 Å [O I] line, which is usually assumed to yield reliable abundances due to its freedom from NLTE departures (Altrock 1968, Lambert 1978). These uncertainties mainly arise from the Solar equivalent width and the method of normalization of stellar O abundances to Solar. Additionally, the simple two-stream calculations of Tomkin *et al.* (1992) indicated that the effects of thermal inhomogeneities on derived O abundances were larger for the [O I] line than for the O I triplet in the case of the cool halo dwarf HD 103095 (Gmb 1830). Thus, while we are *not* saying that O abundances from the 6300 Å line or the 7774 Å triplet are wrong, we do conclude that the *uncertainties* in the 6300 Å abundances may be just as large as for the 7774 Å triplet.

While the 7774 Å O I abundances may be reliable for those Pop I dwarfs having $T \leq 6200\text{--}6300$ K, we must ask if this is also the case for more metal-poor stars. Indeed, Kiselman's (1991) calculations (which might best be regarded with some degree of caution given the inability of these to predict the

solar 7774 Å equivalent widths and the non-inclusion of collisions with neutral H atoms) predict that NLTE effects on the 7774 Å O I abundances increase with decreasing metallicity ($[\text{Fe}/\text{H}]$) and with declining stellar gravity ($\log g$ tends to be lower in the halo dwarfs usually observed than in the Pop I dwarfs analyzed in chapter 5). If one adopts the canonical belief that the 6300 Å [O I] abundances are accurate, then our revised T_{eff} scale (which is consistent with: the results of the new Kurucz 1991 results and the Saxner & Hammarback 1985 relations, a variety of spectroscopic T_{eff} estimates, the Balmer line fitting results of Fuhrmann *et al.* 1993, the Bell & Gustafsson 1989 model calculations, and the color transformations of the Victoria and Yale isochrones) gives 7774 Å O abundances which must also be judged to be accurate. The revised $[\text{O}/\text{Fe}]$ ratios for metal-poor stars (presented in chapter 2) agree with those found for the two most metal-poor dwarfs by Spite & Spite (1991), who used the 6300 Å line. As seen in chapter 4, the mean revised halo dwarf $[\text{O}/\text{Fe}]$ ratios are also in good accord with those derived for giants. A particular concern is that $[\text{O}/\text{Fe}]$ ratios in giants could be affected by O→N cycling and subsequent mixing. In chapter 6, we identified observational clues indicating that present stellar models may be incomplete in that additional mixing seems warranted at evolved stages. Indeed, it has been suggested that such additional mixing may play a part in the large spread of O abundances seen in globular cluster giants. If this is the case, then the relevant quantity of interest may be the *maximum* $[\text{O}/\text{Fe}]$ ratios exhibited by field giants. Using the field giant data of Sneden *et al.* (1991, 1992) and Kraft *et al.* (1992), who have probably performed the highest quality analyses to date, one finds the maximum ratios are $[\text{O}/\text{Fe}] \sim +0.5$, which are in very good agreement with our mean halo dwarf ratio ($[\text{O}/\text{Fe}] = +0.53$). We emphasize again, however, that relatively small effects ($\leq 0.05 - 0.10$ dex or so) on the 7774 Å O abundances cannot be excluded given the uncertainties.

Armed with the above conclusions, we make the following inferences. The scatter seen in the halo dwarf $[\text{O}/\text{Fe}]$ ratios (or the scatter seen in the giant $[\text{O}/\text{Fe}]$ ratios utilized in chapter 4) is completely consistent with the abundance measurement uncertainties (*i.e.*, there is no discernable “cosmic” scatter). This has two implications:

- The SFR in the early Galaxy did not undergo significant variations, *i.e.*, it is doubtful that our Galaxy was a starburst system. Inferences about other spiral galaxies based solely on our own would be overly speculative.
- The early halo was relatively homogeneous, *i.e.*, formation of the Galaxy by mergers of independent “fragments” (*e.g.*, Searle & Zinn 1978), each having their own star formation and chemical evolution history, does not seem to be consistent with the $[\text{O}/\text{Fe}]$ data.

The revised halo dwarf data from chapter 2 and the red giant data from chapter 4 both suggest that $[\text{O}/\text{Fe}]$ is constant in the halo, at least for $[\text{Fe}/\text{H}] \geq -3$. Now, the change in behavior of $[\text{O}/\text{Fe}]$ in going from such low $[\text{Fe}/\text{H}]$ values to Solar values (as seen in Figure 16) can be interpreted, as is well-known, as O and Fe being produced in different sites (Type II and Type I supernovae, respectively) having different evolutionary timescales (with the implicit assumption that $[\text{Fe}/\text{H}]$ grossly serves as some sort of chronometer). Given the constancy and this evolutionary timescale argument:

- The duration of halo formation is set by the timescale of Type Ia supernovae production of Fe. Calculations assuming that CO–CO white dwarfs are the dominant source of Type Ia supernovae indicate that this timescale is a $\leq 10^9$ yr (Smecker-Hane & Wyse 1992). However, age spreads of $\sim 3 \times 10^9$ yr inferred for both globular clusters and halo field stars (*e.g.*, Schuster & Nissen 1989) are inconsistent with this theoretical time scale. Recent work, though, provides a possible way for this inconsistency to be overcome:

- The apparent inconsistency between age spreads in the halo and current Type Ia timescales can be resolved if CO–He white dwarfs are the dominant Type Ia sources as suggested by Smecker-Hane & Wyse (1992). These sources have Fe production time scales of order $10^9 - 10^{10}$ yr.

As these authors note, there are problems with such a proposal. Besides the theoretical difficulties encountered (*e.g.*, the sensitivity of the calculations to the assumed mass transfer rate or the fact that some calculations indicate that CO–He systems may not go supernovae), the CO–He timescale may be *too* long such that type II systems enrich the disk gas above observed levels (Gilmore *et al.* 1989, Smecker-Hane & Wyse 1992). However, this difficulty might be overcome if there is a substantial delay between the end of halo formation and the beginning of star formation in the disk. Is there evidence for this? Wheeler *et al.* (1989) qualitatively demonstrate how such a delay can lead to gaps in the observed $[\text{O}/\text{H}]$ distribution. Therefore, it is important to note the following:

- Wheeler *et al.* (1989) suggested that a gap may be present in the $[\text{O}/\text{H}]$ distribution in the approximate region $-0.8 \lesssim [\text{O}/\text{H}] \lesssim -0.5$. In chapter 4, it was shown that this gap apparently persists despite the additional data. Such a gap may be evidence of a hiatus between the end of halo formation and the beginning of star formation in the disk. This hiatus would alleviate the problems caused by such a long CO–He Type Ia supernova timescale.

- Furthermore, the open cluster results presented in chapter 6 are consistent with Type Ia

supernovae time scales as long as a few Gyr. It was seen there that a clear $[\text{O}/\text{H}] - \text{Age}$ relation was present over ~ 7 Gyr, as might be expected since the Type II supernova (which are the predominant sources of O) timescale is very short (10^8 yr). For these same clusters, however, no relation was seen between $[\text{Fe}/\text{H}]$ and age. This might be somewhat unexpected if the timescale for Type Ia supernovae (the predominant sources of Fe) is truly $\leq 10^9$ yr. However, such a circumstance might be expected if the Type Ia timescale is significantly longer (few Gyr) as predicted for CO-He white dwarf systems. Such a long timescale for Fe production may also produce the significant scatter seen in the $[\text{Fe}/\text{H}] - \text{Age}$ relations of, *e.g.*, Edvardsson *et al.* (1993).

In fairness, it must be readily acknowledged that such arguments are not foolproof. One can not exclude fragment merger models at present. Such models are desirable because since they also allow slow formation of the halo, thus accommodating age spreads among globular clusters and among field stars. Indeed, we note that some of the O data would actually be consistent with such hierarchical merger models:

- In halo formation models which propose merging of independent fragments, the break in the $[\text{O}/\text{Fe}]$ *vs.* $[\text{Fe}/\text{H}]$ plane is expected to be poorly defined. This is because chemical evolution is expected to proceed independently in each fragment. We noted in chapter 4 that simple regressions do not favor any unique value of the $[\text{O}/\text{Fe}]$ break point. Thus, it may well be (as simple inspection of Figure 16 might indicate) that the break is not well defined.

- The independent fragments in a merger model must evolve as such to produce the observed mean Fe abundance ($[\text{Fe}/\text{H}] \sim -1.5$) of the halo. Therefore, not many fragments can have a metallicity of $[\text{Fe}/\text{H}] \sim -1.0$. If $[\text{O}/\text{Fe}]$ is constant in the halo at $+0.5$, then this simple fact alone may result in a shortfall of stars having $-0.8 \lesssim [\text{O}/\text{H}] \lesssim -0.5$ as is apparently observed.

Clearly, much work remains to be done and the nature of the Galaxy's formation still remains elusive and esoteric. In particular, the large observed scatter in r process elements such as Eu may indicate inhomogeneity in the halo formation process and favor mergers of independent fragments.

2. Other Conclusions

Several other conclusions are also reached. Possible cosmological implications of the suggested higher T_{eff} scale for metal-poor stars were considered. Important, but somewhat unexciting, inferences were drawn. First of all, the B/Be ratio of HD 140283 was raised slightly from 10 to 12. The

latter value is more comfortably consistent with ordinary spallation production of B and Be and not with significant primordial abundances of these elements.

Secondly, the higher T_{eff} scale seems to be consistent with both the Victoria and the Yale groups' isochrone color- T_{eff} transformations. Thus, these are seemingly eliminated as a significant source of uncertainty in globular cluster ages. The issue of the effects of the [O/Fe] ratio itself on globular cluster ages cannot be addressed until agreement can be reached concerning the age effects of raising the α element abundances in the models. If suggestions (*e.g.*, Chieffi, Straniero & Salaris 1991) that including overabundances of the α elements (Mg, Ca, *etc.*) as well as O does not result in a reduction of globular cluster ages is correct, then the ages yielded by the very newest Victoria and Yale isochrones may need to be *raised* (by ~ 2 Gyr or so). Even if such suggestions are wrong, ages yielded by the Victoria isochrones for the most metal-poor globular clusters, which are assumed to have [O/Fe] ratios larger than present observations indicate, may need to be raised slightly (by $\lesssim 1$ Gyr).

The third cosmological issue examined was the primordial Li abundance. These results, presented in chapter 3, have already been superseded by the newer data and analyses of Thorburn (1994). In what is likely to become the seminal work on halo star Li abundances for some time to come, she has found that the mean halo Li abundance is $N(\text{Li}) \sim 2.3$, substantially larger than the value of ~ 2.1 determined previously by others. If one follows the liberal procedure of assuming that Thorburn's value of $N(\text{Li}) \sim 2.3$ represents the true level of the Spite plateau, then applies a correction for depletion as calculated from standard models (as done in chapter 3), and further adjusts her Li abundances upward by 0.1 dex for consistency with the T_{eff} scale advocated here, then one arrives at a value of $N(\text{Li}) \sim 2.5$ for the primordial Li abundance. While we emphasize that this was *not* the conclusion of Thorburn (1994, who suggests that Li enrichment has occurred and that the genuine primordial Li abundance is $N(\text{Li}) \sim 2.2$), we note that even the more liberal approach (which includes our revised T_{eff} scale) yields a value ($N(\text{Li}) \sim 2.5$) which is still consistent with standard models of Big Bang nucleosynthesis given all the uncertainties. However, as observed abundances go higher, it is all that much easier for rotational stellar models (*e.g.*, Pinsonneault, Deliyannis & Demarque 1992) to deplete a very high primordial Li abundance ($N(\text{Li}) \sim 3.0$, inconsistent with standard Big Bang models) to the observed level.

Finally, using the 7774 Å O I triplet, we derived O abundances for a small sample of open clusters having a range of ages. Poor agreement is found between our Hyades and UMa group abundances and those of Garcia Lopez *et al.* (1993). Given the good agreement we found between the Pop I field star O abundances from the 7774 Å O I triplet and the 6300 Å [O I] line, this

disagreement underscores the suspect nature of present NLTE O calculations. As noted above, the clusters exhibit a clear relation between age and O abundance which contrasts with the lack of any correlation between age and Fe abundance. It is suggested that O abundances may be a superior chronometer in studying the chemical evolution of the Galactic disk. This certainly would be true if the Type Ia supernovae, responsible for most Fe production, have timescales of several Gyr as suggested above. In any case, it is clear that the $[\text{Fe}/\text{O}]$ vs. $[\text{O}/\text{H}]$ plane offers a much different and better perspective in studying the chemical evolution of the open cluster system than the traditional $[\text{O}/\text{Fe}]$ vs. $[\text{Fe}/\text{H}]$ plane; this seemingly verifies the suggestion made by Wheeler *et al.* (1989).

The O abundances we see in the young clusters are supersolar. This exacerbates the disagreement between O abundances derived for the Sun and solar-type stars and those abundances presumed to represent the present-day ISM as derived from planetary nebulae, H II regions, and supergiants. It is suggested that the planetary nebulae and supergiant O abundances may be severely affected by additional mixing not included in present models. This is supported by several pieces of observational evidence. If one considers this additional mixing, then it was shown in chapter 6 that the planetary nebulae and most recent supergiant abundances are entirely consistent with the cluster O abundances. It is further suggested that the O abundances derived from gaseous nebulae may be affected by O being locked up in dust grains. If this resolution to the present-day O abundance discrepancy is correct, then inferences made about extragalactic O abundances or Galactic gradients in O abundance may be incorrect.

3. Future Work

The work presented here has, as normal, raised many more questions than it has answered. The above summary of our conclusions specifically identifies numerous pieces of work to be performed. Hydrodynamical calculations should be carried out using metal-poor models in order to calculate the magnitude of possible thermal inhomogeneities in the atmospheres of halo stars. Abundance analyses incorporating these inhomogeneities should be performed in order to determine the effects of O abundances derived from *all* O spectral features. NLTE calculations clearly deserve continued attention as they are not even sufficiently developed to match Solar observations. The hotter T_{eff} values for metal-poor stars advocated in this dissertation need to be confirmed. This can be done by performing more halo star fine analyses like those of Smith *et al.* (1992) and Fuhrmann *et al.* (1993). If these higher T_{eff} values are correct, then the redetermination of ratios such as $[\text{Eu}/\text{Fe}]$ is clearly warranted in order to see if the large scatter presently observed, perhaps suggesting an

inhomogeneous halo formation, persists. Fragmentation models seem to require that a substantial fraction of gas is lost from the halo during collapse (Gilmore *et al.* 1989). Since this gas has been proposed to flow to the Galactic center and form the bulge, it would be of obvious interest to determine [O/Fe] ratios in bulge stars to see if they are consistent with ratios observed in the halo. This may be possible for *in situ* samples of bulge giants using the new generation of large aperture telescopes. Alternatively, one might hope to identify potential bulge stars passing through the solar neighborhood on the basis of their kinematics.

Further investigation of the Yale rotational models is urgently needed. The effects of rotational mixing on the observed O abundances in halo field and globular cluster giants as well as disk planetary nebulae and supergiant progenitors must be established in order to assess if their observed O abundances truly represent their initial O abundances. Since the Kurucz (1991) models and the Bell & Gustafsson (1989) models seem to be consistent with our revised T_{eff} values, it would seem important to investigate the effects of using these atmospheres to establish the boundary conditions of the Yale rotation models.

Given the encouraging results of the open cluster O abundances, much future work comes to mind. First, observations should be obtained of younger and older clusters. Several good candidates (α Per, Mel 66, and NGC 6791) are available. However, determinations for the most distant clusters may require use of the 6300 Å [O I] line in order to avoid the limitations of OH sky emission on the 7774 Å abundances. More distant clusters and anti-center clusters must also be observed. Only then can issues such as gradients be addressed. It would also seem worthwhile to examine the question of Age – metallicity relations in disk field stars using O as a metallicity indicator. Such an attempt might use a carefully constructed sample such as that of Twarog (1980) and also derive spectroscopic [Fe/H] values. A definite advantage of a spectroscopic approach is the ability to exclude problematic objects such as Fm stars. Above, we suggested that O abundances of gaseous nebulae may be affected by O being locked up in dust grains. As noted in chapter 6, the usual argument against this is the relative constancy of derived Ne-O ratios. Future work must examine whether this constancy holds given the possibility of temperature fluctuations along the line of sight to such nebulae. The author hopes that this dissertation will encourage such future work as it will be needed to improve our understanding of the chemical evolution of oxygen in the Galaxy.

Literature Cited

- Abia, C. & Rebolo, R. 1989, ApJ, 347, 186 (AR)
- Altrock, R. 1968, Sol. Phys., 5, 260
- Andersen, J., Edvardsson, B., Gustafsson, B. & Nissen, P. E. 1988, in *The Impact of Very High S/N Spectroscopy on Stellar Physics*, IAU Symposium 132, ed. G. Cayrel de Strobel and M. Spite, (Dordrecht: Kluwer), p. 441
- Artjukhina, N. M. 1966, Tr. Gos. Astron. Inst. Shternberga, 35
- Artjukhina, N. M. 1971, Soobshch. Gos. Astron. Inst. Shternberga, 172, 3
- Balachandran, S. 1990, ApJ, 354, 310
- Balachandran, S., Carney, B. W., Fry, A. M., Fullton, L. K. & Peterson, R. C. 1993, ApJ, 413, 368
- Barbuy, B. 1988, A&A, 191, 121
- Barbuy, B. & Erdelyi-Mendes, M. 1989, A&A, 214, 239
- Barbuy, B. 1990, in *Cool Stars, Stellar Systems and the Sun*, edited by G. Wallerstein, (San Francisco: ASP), p.328
- Barry, D. C. 1988, ApJ, 334, 436
- Baschek, B., Scholz, M., Sedlmayer, E. 1977, A&A, 55, 375
- Beers, T. C., Preston, G. W., Shectman, S. A. & Kage, J. A. 1990, AJ, 100, 849
- Bell, R. A. & Gustafsson, B. 1989, MNRAS, 236, 653
- Bergbusch, P. A. & Vandenberg, D. A. 1992, ApJS, 81, 163 (BV92)
- Bessell, M. S., Sutherland, R. S. & Ruan, K. 1991, ApJ, 383, L71
- Bevington, P. R. 1969, *Data Reduction and Error Analysis for the Physical Sciences*, (New York: McGraw-Hill Inc.), p. 195
- Biemont, E., Baudoux, M., Kurucz, R. L., Ansbacher, W. & Pinnington, E. H. 1991, A&A, 249, 539
- Biemont, E., Hibbert, A., Godefroid, M., Vaeck, N. & Fawcett, B. C. 1991, ApJ, 375, 818
- Boesgaard, A. M. & Steigman, G. 1985, ARA&A, 23, 319
- Boesgaard, A. M. & Tripicco, M. J. 1987, ApJ, 313, 389
- Boesgaard, A. M. & Budge, K. G. 1988, ApJ, 332, 410
- Boesgaard, A. M., Budge, K. G. & Burck, E. E. 1988, ApJ, 325, 749
- Boesgaard, A. M., Budge, K. G. & Ramsay, M. K. 1988, ApJ, 327, 389
- Boesgaard, A. M. 1989, ApJ, 336, 798
- Boesgaard, A. M. & Friel, E. D. 1990, ApJ, 351, 467
- Boesgaard, A. M. & King, J. R. 1993, AJ, in press (December 1993 volume)

- Breger, M. 1986, ApJ, 309, 311
- Brown, J. A. 1985, ApJ, 297, 233
- Brown, T. M., Demarque, P., Noyes, R., Praderie, F., Roxburgh, I. W. & Schatzman, E. 1993, ASP Conf. Ser., 40, 776
- Carlberg, R. G., Dawson, P. C., Hsu, T. & Vandenberg, D. A. 1985, ApJ, 294, 674
- Carney, B. W. 1983, AJ, 88, 623
- Carney, B. W. 1992, private communication
- Caputo, F., Chieffi, A., Castellani, V., Collados, M., Martinez-Roger, C. & Paez, E. 1990, AJ, 99, 261
- Cayrel de Strobel, G. & Bentolila, C. 1983, A&A, 119, 1
- Cayrel, R., Cayrel de Strobel, G. & Campbell, B. 1985, A&A, 146, 249
- Cayrel, R. 1988, in *The Impact of Very High S/N Spectroscopy on Stellar Physics*, eds. G. Cayrel de Strobel and M. Spite, (Dordrecht: Kluwer), p. 345
- Cayrel de Strobel, G., Hauck, B., Thevenin, F., Friel, E. Mermilliod, M. & Borde, S. 1992, A&AS, 95, 273
- Chaboyer, B., Deliyannis, C. P., Demarque, P., Pinsonneault, M. H. & Sarajedini, A. 1992, ApJ, 388, 372
- Chieffi, A., Straniero, O. & Salaris, M. 1991, in *The Formation and Evolution of Star Clusters*, ed. K. Janes, ASP Conf. Ser., 13, 219
- Clegg, R. E. S., Lambert, D. L. & Tomkin, J. 1981, ApJ, 250, 262
- Cohen, J. G., Frogel, J. A. & Persson, S. E. 1978, ApJ, 222, 165 (CFP)
- Cohen, J. G. 1979, ApJ, 231, 751
- Cohen, J. G. 1980, ApJ, 241, 981
- Cohen, J. G. 1983, ApJ, 270, 654
- Conti, P. S., Greenstein, J. L., Spinrad, H., Wallerstein, G. & Vardya, M. S. 1967, ApJ, 148, 105
- Crawford, D. L. & Barnes, J. V. 1969a, AJ, 74, 407
- Crawford, D. L. & Barnes, J. V. 1969b, AJ, 74, 818
- Crawford, D. L. & Barnes, J. V. 1970, AJ, 75, 946
- Crawford, D. L. & Perry, C. L. 1976, AJ, 81, 419
- de Freitas Pacheco, J. A. 1993, ApJ, 403, 673
- Deliyannis, C. P., Demarque, P., Kawaler, S. D., Krauss, L. M. & Romanelli, P. 1989, Phys. Rev. Lett., 62, 1583
- Deliyannis, C. P., Demarque, P. & Kawaler, S. D. 1990, ApJS, 73, 21

- Deliyannis, C. P., Pinsonneault, M. H. & Duncan, D. K. 1993, *ApJ*, 414, 740
- Demarque, P., Green, E. M. & Guenther, D. B. 1992, *AJ*, 103, 151
- Duncan, D. K., Lambert, D. L. & Lemke, M. 1992, *ApJ*, 401, 584
- Edmunds, M. G., Greenhow, R. M., Johnson, D., Kluckers, V. & Vila, M. B. 1991, *MNRAS*, 251, 33p
- Edvardsson, B., Andersen, J., Gustafsson, B., Lambert, D. L., Nissen, P. E. & Tomkin, J. 1993, *A&A*, 275, 101
- Eggen, O. J. 1963, *ApJ*, 138, 356
- Eggen, O. J. & Sandage, A. R. 1964, *ApJ*, 140, 130
- Eggen, O. J. 1969, *PASP*, 81, 553
- Eggen, O. J. & Sandage, A. 1969, *ApJ*, 158, 669
- Eggen, O. J. 1971, *ApJS*, 22, 389
- Eggen, O. J. 1978, *ApJS*, 37, 251
- Feigelson, E. D. & Babu, G. J. 1992, *ApJ*, 397, 55 (FB)
- Francic, S. P. 1989, *AJ*, 98, 888
- Friel, E. D. & Boesgaard, A. M. 1990, *ApJ*, 480
- Friel, E. D. & Boesgaard, A. M. 1992, *ApJ*, 387, 170
- Friel, E. D. & Janes, K. A. 1993, *A&A*, 267, 75
- Frogel, J. A., Persson, S. E. & Cohen, J. G. 1981, *ApJ*, 246, 842
- Frogel, J. A., Persson, S. E. & Cohen, J. G. 1983, *ApJS*, 53, 713
- Fuhrmann, K., Axer, M. & Gehren, T. 1993, *A&A*, 271, 451
- Fuller, G. M., Boyd, R. N. & Kalen, J. D. 1991, *ApJ*, 371, L11
- Garcia Lopez, R. J., Rebolo, R., Herrero, A. & Beckman, J. E. 1993, *ApJ*, 412, 173
- Geisler, D., Claria, J. J. & Minniti, D. 1992, *AJ*, 104, 1892
- Gilmore, G., Wyse, R. F. G. & Kuijken, K. 1989, *ARA&A*, 27, 555
- Gilmore, G., Gustafsson, B., Edvardsson, B. & Nissen, P. E. 1992, *Nature*, 357, 379
- Girard, T. M., Grundy, W. M., Lopez, C. E. & van Altena, W. F. 1989, *AJ*, 98, 227
- Gratton, R. G. & Ortolani, S. 1986, *A&A*, 169, 201
- Gratton, R. G. & Sneden, C. 1987, *A&A*, 178, 179
- Gratton, R. G. & Sneden, C. 1991, *A&A*, 241, 501
- Gray, D. F. 1992, *The Observation and Analysis of Stellar Photospheres*, 2nd edition, (Cambridge University Press, Cambridge)

- Green, E. M., Demarque, P. & King, C. R. 1987, *The Revised Yale Isochrones and Luminosity Functions*, New Haven: Yale University Observatory
- Green, E. M. 1988, in *Calibration of Stellar Ages*, ed. A. G. Davis Philip, (Schenectady: L. Davis Press), p. 81 (G88)
- Grevesse, N. & Anders, E. 1989, in *Cosmic Abundances of Matter*, ed. C. J. Waddington, (New York: AIP), p. 1
- Griffin, R. F., Gunn, J. E., Zimmerman, B. A., Griffin, R. E. M. 1988, AJ, 96, 1988
- Gurtovenko, E. A. & Kostik, R. I. 1981, A&AS, 46, 239
- Gustafsson, B., Bell, R. A., Eriksson, K. & Nordlund, A. 1975, A&A, 42, 407
- Hauck, B. & Mermilliod, M. 1990, A&AS, 86, 107
- Hayes, D. S. & Latham, D. W. 1975, ApJ, 197, 593
- Heinemann, K. 1926, AN, 227, 193
- Hobbs, L. M., Thorburn, J. A. & Rodriguez-Bell, T. 1990, AJ, 100, 710
- Hobbs, L. M. & Thorburn, J. A. 1991, ApJ, 375, 116
- Hobbs, L. M. & Thorburn, J. A. 1991, AJ, 102, 1070
- Hobbs, L. M. & Thorburn, J. A. 1992, AJ, 104, 669
- Hoffleit, D. & Jaschek, C. 1982, *The Bright Star Catalog*, 4th Revised Edition, (New Haven: Yale University Observatory)
- Holweger, H., Heise, C. & Kock, M. 1990, A&A, 232, 510
- Holweger, H., Bard, A., Kock, A. & Kock, M. 1991, A&A, 249, 545
- Isobe, T., Feigelson, E. D., Akritas, M. G. & Babu, G. J. 1990, ApJ, 364, 104 (IFAB)
- Janes, K. A. 1979, ApJS, 39, 135
- Johnson, D. R. H. & Soderblom, D. R. 1987, AJ, 93, 864
- Johnson, H. L. 1952, ApJ, 116, 640
- Johnson, H. L. 1953, ApJ, 117, 357
- Johnson, H. L. & Knuckles, C. 1955, ApJ, 122, 209
- Johnson, H. L. & Mitchell, R. I. 1958, ApJ, 128, 31
- Jones, B. F. & Cudworth, K. 1983, AJ, 88, 215
- Jones, B. F. & Stauffer, J. R. 1991, AJ, 102, 1082
- Kawano, L. H., Fowler, W. A., Kavanagh, R. W. & Malaney, R. A. 1991, ApJ, 372, 1
- King, J. R. 1993a, AJ, 106, 1206
- King, J. R. 1993b, AJ, in press (January 1994 volume)
- Kiselman, D. 1991, A&A, 245, L9

- Kjaergaard, P., Gustafsson, B., Walker, G. A. H. & Hultqvist, L. 1992, A&A, 115, 145
- Klein-Wassink, W. J. 1927, Publ. Kapteyn Astron. Lab., 41
- Kraft, R. P., Sneden, C., Langer, G. E., Prosser, C. F. 1992, AJ, 104, 645
- Krauss, L. M. & Romanelli, P. 1990, ApJ, 358, 47
- Kurucz, R. L. 1979, ApJS, 40, 1
- Kurucz, R. L., Furenlid, I., Brault, J. & Testerman, L. 1984, Solar Flux Atlas from 296 to 1300 nm,
National Solar Observatory Atlas No. 1 (KFBT, K84)
- Kurucz, R. L. 1991, CfA Preprint No. 3348
- Kurucz, R. L. 1992, private communication
- Laird, J. B. 1985, ApJ, 289, 556
- Laird, J. B., Carney, B. W. & Latham, D. W. 1988, AJ, 95, 1843
- Lambert, D. L. 1978, MNRAS, 182, 249
- Lambert, D. L. 1989, in *Cosmic Abundances of Matter*, ed. C. J. Waddington, (New York: AIP),
p. 168
- Lambert, D. L., Heath, J. E. & Edvardsson, B. 1991, MNRAS, 253, 610
- Luck, R. E. & Bond, H. E. 1981, ApJ, 244, 919
- Luck, R. E. & Bond, H. E. 1985, ApJ, 292, 559
- Maeder, A. & Meynet, G. 1988, A&AS, 76, 411
- Magain, P. 1987a, A&A, 181, 323 (M87)
- Magain, P. 1987b, A&A, 179, 176
- Magain, P. 1989, A&A, 209, 211
- Mathieu, R. D., Latham, D. W., Griffin, R. F. & Gunn, J. E. 1986, AJ, 92, 1100
- Matteucci, F. & Francois, P. 1992, A&A, 262, L1
- Meusinger, H. 1992a, ApSS, 188, 19
- Meusinger, H. 1992b, A&A, 266, 190
- Montgomery, K. A., Marschall, L. A. & Janes, K. A. 1993, AJ, 106, 181
- Moore, C. E., Minnaert, M. G. J. & Houtgast, J. 1966, NBS Monograph 61
- Nissen, P. E. 1981, A&A, 97, 45
- Nissen, P. E., Twarog, B. A. & Crawford, D. L. 1987, AJ, 93, 635
- Nissen, P. E. 1988, A&A, 199, 146
- Nissen, P. E. & Edvardsson, B. 1992, A&A, 261, 255 (NE92)
- Nordlund, A. & Dravins, D. 1990, A&A, 228, 155
- Osterbrock, D. E., Tran, H. D. & Veilleux, S. 1992, ApJ, 389, 305

- Peimbert, M. Storey, P. J. & Torres-Peimbert, S. 1993, ApJ, 414, 626
- Penn, M. J., Mickey, D. L., Canfield, R. C. & LaBonte, B. J. 1991, Sol Phys, 135, 163
- Perrin, M. N., Hejlesen, P. M., Cayrel de Strobel, G. & Cayrel, R. 1977, A&A, 54, 779
- Pinsonneault, M. H., Deliyannis, C. & Demarque, P. 1992, ApJS, 78, 179
- Platais, I. 1991, A&AS, 87, 69
- Prantzos, N., Casse, M. & Vangioni-Flam, E. 1993, ApJ, 403, 630
- Rana, N. C. & Basu, S. 1992, A&A, 265, 499
- Read, S. M. & Viola, V. E. Jr. 1984, Atomic Data. Nucl. Data Tables, 31, 359
- Rebolo, R., Molaro, P. & Beckman, J. E. 1988, A&A, 192, 192
- Reglero, V. & Fabregat, J. 1991, A&AS, 90, 25
- Renzini, A. & Voli, M. 1981, A&A, 94, 175
- Ridgway, S. T., Joyce, R. R., White, N. M. & Wing, R. F. 1980, ApJ, 235, 126
- Ridgway, S. T., Jacoby, G. H., Joyce, R. R., Siegel, M. J. & Wells, D. C. 1984, AJ, 87, 1044
- Ryan, S. G. & Norris, J. E. 1991a, AJ, 101, 1865
- Ryan, S. G. & Norris, J. E. 1991b, AJ, 101, 1835
- Ryan, S. G., Norris, J. E., Bessell, M. S. & Deliyannis, C. P. 1992, ApJ, 388, 184
- Sandage, A. R. 1962, ApJ, 135, 33
- Sandage, A. 1988, in Calibration of Stellar Ages, ed. A.G. Davis Philip, (Schenectady: L. Davis Press), p. 43
- Sanders, W. L. 1977, A&AS, 27, 89
- Saxner, M. & Hammarbäck, G. 1985, A&A, 151, 372 (SH)
- Schuster, W. J. & Nissen, P. E. 1988, A&AS, 73, 225
- Schuster, W. J. & Nissen, P. E. 1989, A&A, 222, 69
- Searle, L. & Zinn, R. 1978, ApJ, 225, 357
- Smecker-Hane, T. A. & Wyse, R. F. G. 1992, AJ, 103, 1621
- Smith, G., Lambert, D. L. & Ruck, M. J. 1992, A&A, 263, 249
- Snedden, C., Lambert, D. L. & Whitaker, R. W. 1979, ApJ, 234, 964
- Snedden, C., Kraft, R. P., Prosser, C. F. & Langer, G. E. 1991, AJ, 102, 2001
- Snedden, C., Kraft, R. P., Prosser, C. F. & Langer, G. E. 1992, AJ, 104, 2121
- Sofia, S., Pinsonneault, M., & Deliyannis, C. P. 1991, in Angular Momentum Evolution of Young Stars, edited by S. Catalano and J.R. Stauffer, (Kluwer, Dordrecht), p. 333
- Soderblom, D. R. & Mayor, M. 1993, AJ, 105, 226
- Spiesman, W. J. & Wallerstein, G. 1991, AJ, 102, 1790

- Spite, F. & Spite, M. 1982, A&A, 115, 357
- Spite, F. & Spite, M. 1986, A&A, 163, 140
- Spite, M. & Spite, F. 1991, A&A, 252, 689
- Stetson, P. B. & Harris, W. E. 1988, AJ, 96, 909
- Stothers, R. B. & Chin, C.-W. 1992, ApJ, 390, 136
- Thorburn, J. A. & Beers, T. C. 1993, ApJ, 404, L13
- Thorburn, J. A. 1994, ApJ, in press (January 10 volume)
- Tomkin, J., Sneden, C. & Lambert, D. L. 1986, ApJ, 302, 415
- Tomkin, J., Lemke, M., Lambert, D. L. & Sneden, C. 1992, AJ, 104, 1568 (TLLS)
- Twarog, B. A. 1980, ApJ, 242, 242
- Twarog, B. A. & Wheeler, J. C. 1982, ApJ, 261, 636
- Twarog, B. A. 1983, ApJ, 267, 207
- Twarog, B. A. & Anthony-Twarog, B. J. 1989, AJ, 97, 759
- Unsold, A. 1955, Physik der Sternatmosphären, 2nd edition, (Springer-Verlag, Berlin)
- Ugoren, A. R., Mesrobian, W. S. & Kerridge, S. J. 1972, AJ, 77, 74
- VandenBerg, D. A. & Bell, R. A. 1985, ApJS, 58, 561
- VandenBerg, D. A. 1988, in *The Extragalactic Distance Scale*, ed. S. van den Bergh and C. J. Pritchett, (Provo: BYU Print Services), p. 187
- VandenBerg, D. A. & Poll, H. E. 1989, AJ, 98, 1451
- VandenBerg, D. A. & Stetson, P. B. 1991, AJ, 102, 1043
- VandenBerg, D. A. 1992, ApJ, 391, 685 (V92)
- Venn, K. A. 1993, ApJ, 414, 316
- Walter, D. K., Dufour, R. J. & Hester, J. J. 1992, ApJ, 397, 196
- Wheeler, J. C., Sneden, C. & Truran, J. W., Jr. 1989, ARA&A, 27, 279
- Wiese, W. L., Smith, M. W. & Glennon, B. M. 1966, NSR DS-NBS 4
- Wyse, R. F. G. & Gilmore, G. 1992, AJ, 104, 144