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THE CHEMICAL COMPOSITION OF THREE LAMBDA BOOTIS STARS

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ABSTRACT

Abundance analyses are reported for three certain members (λ Boo, 29 Cyg, π^1 Ori) of the class of rapidly rotating, metal-poor A-type stars known as λ Boötis stars. Model atmosphere analysis of high-resolution, high signal-to-noise spectra shows that the metal deficiencies are more severe than previously reported: [Fe/H] = −2.0, −1.8, −1.3 for λ Boo, 29 Cyg, and π^1 Ori, respectively. Other metals (Mg, Ca, Ti, and Sr) are similarly underabundant with Na often having a smaller underabundance. C, N, O, and S have near-solar abundances. Vega is shown to be a mild λ Boo star.

The abundance anomalies of the λ Boo stars resemble those found for the interstellar gas in which the metals are depleted through formation of interstellar grains. It is suggested that the λ Boo stars are created when circumstellar (or interstellar) gas is separated from the grains and accreted by the star. The bulk of the interstellar grains comprises a circumstellar cloud or disk that is detectable by its infrared radiation.

Subject headings: stars: abundances — stars: accretion — stars: peculiar A

I. INTRODUCTION

The main sequence around spectral type A is home for a bewildering variety of stars exhibiting anomalously weak or strong absorption lines with respect to "normal" A-type stars for which Vega is commonly taken as representative. Today, the stars with anomalous spectra are frequently referred to as chemically peculiar (CP) stars. This convenient label encompasses the common Ap(Si), Ap(HgMn), Ap(SrCrEu), and the Am stars as well as a few rarer groups, The domain of the CP stars spreads into the adjacent spectral types (B and F). Moreover, many or even all normal A stars show abundance anomalies. Holweger, Steffen, and Gigas (1986) preface a discussion of high-resolution spectra with an opposite remark by W. P. Bidelman, a pundit of spectral classification and spectral peculiarities—"I do not know, nor have I ever seen, a normal A0 star" (see Eggen 1984). Recent work by Holweger, Gigas, and Steffen (1986) and Holweger, Steffen, and Gigas (1986) has demonstrated that early A stars classified as normal show striking abundance anomalies.

In this paper, we present an abundance analysis for members of an almost forgotten class of unusual A stars: the λ Bootis stars. Discovered by Morgan, Keenan, and Kellman (1943), λ Boo with a spectral type near A0 was noted to have very weak lines of the common metals such as Mg and Ca. Recently, several authors have rediscussed the definition of the class and suggested potential candidates for membership (e.g., Hauck and Slettebak 1983; Abt 1984; Gray 1988). Our goal was to derive the chemical composition of three stars universally accepted as members of the class: the trio includes the eponymous star (λ Boo), π^1 Ori, and 29 Cyg. The initial goal was to obtain accurate abundances for C, N, and O and a selection of heavier elements from Na to Sr. It was expected that the compositions would provide clues to the stars' close relatives and serve to sharpen the definition of a λ Boo star.

Abundance analyses of λ Boo stars have been reported previously. Burbidge and Burbidge (1956) and Baschek and Searle (1969) found that, as was suspected from the work of Morgan, Keenan, and Kellman (1943), the abundances of Mg, Ca, and Fe were substantially less in λ Boo stars than in normal A stars such as Vega. An indication that the underabundance did not extend to oxygen was provided by Kodaira (1967) and Baschek and Searle (1969). Recently, interest in the λ Boo stars has revived. Lambert, McKinley, and Roby (1986) from observations of the C I 9100 Å multiplet noted that carbon was not underabundant. Baschek *et al.* (1984) and Baschek and Slettebak (1988) showed from an examination of ultraviolet (*IUE*) spectra that the abundances of carbon, nitrogen, and oxygen in λ Boo stars were approximately solar, but heavier elements (Mg, Al, Si, S, Mn, Fe, and Ni) were, as found from optical spectra, underabundant by about a factor of 3 relative to normal A stars.

II. OBSERVATIONS AND ANALYSIS

Three λ Boo stars and the "normal" A star Vega were observed—Table 1 lists some basic data for the former stars. A majority of the spectra of the four stars was obtained using the Reticon camera at the coudé focus of the 2.7 m telescope at the McDonald Observatory (Vogt, Tull, and Kelton 1978). For the λ Boo stars, each of the approximately 13 spectral intervals covers 100 Å at a resolution of about 0.25 Å at a signal-to-noise ratio of 100 to 500 per resolution element. A CCD replaced the Reticon for observations of S I lines near 6750, 9220, and 10456 Å. In some intervals, telluric contamination is substantial. To remove the telluric lines, a spectrum of a very rapidly rotating, hotter star was taken at the same air mass. Dividing the spectrum of the λ Boo star by that of the hot star removed the blending telluric lines.

In the case of the λ Boo stars where the absorption lines are broad and often shallow, the accuracy of the equivalent widths is set in large part by the reliability of the flat-fielding technique and the placement of the continuum. Several tests indicate that the equivalent widths are not subject to large systematic errors. The flat-fielded spectra show a slowly sloping and almost linear continuum where the slope is due to the fact that the flat-field lamp has a lower color temperature than the star. In such spectra, the only lines are those present in the sharp-lined spectrum of Vega, and the continuum regions for Vega are continuum regions in the λ Boo stars. When a hot star was observed for the removal of the telluric lines, we compared the flat-fielded spectrum obtained using the flat-field lamp with

TABLE 1

BASIC CHARACTERISTICS OF THE OBSERVED STARS

Quantity[a]	π^1 Ori	λ Boo	29 Cyg
Spectral type	A2 V	A0 V	A0 p
V	4.97	4.65	4.18
$B-V$	0.14	0.09	0.08
$b-y$[b]	0.044	0.051	0.101
m_1	0.178	0.182	0.157
c_1	1.007	1.000	0.927
β	2.898	2.894	2.833
$v \sin i$ (km s^{-1})	110	100	85
π ('' arc)	0.037	0.017	0.033
U (km s^{-1})[c]	−6	−35	−21
V (km s^{-1})	−24	−7	−12
W (km s^{-1})	−10	−5	−3

[a] Unless otherwise indicated entries are taken from the *Bright Star Catalogue* (Hoffleit 1982).

[b] Strömgren photometric indices are from Hauck and Mermilliod 1980.

[c] Velocities are from Hauck and Slettebak 1983.

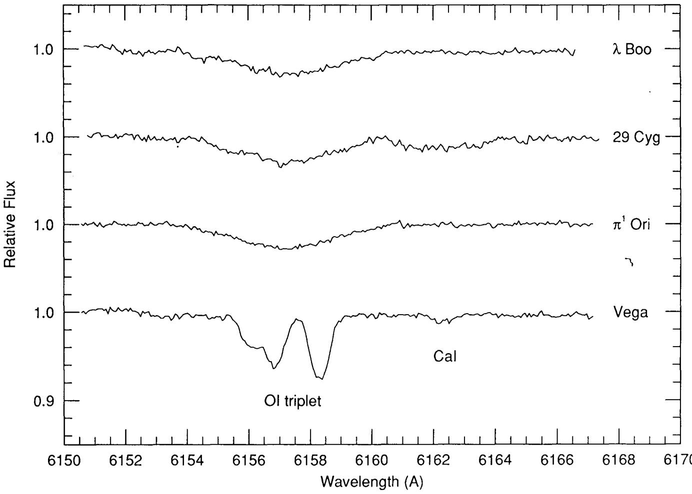

FIG. 2.—Spectra of λ Boo, 29 Cyg, π^1 Ori, and Vega near 6160 Å showing the O I triplet. Spectra of the former three stars have been displaced for clarity.

that obtained using the hot star. The two continua are about equally smooth; spurious lines are absent from both spectra. To within the measurement errors, the equivalent widths are the same for the two spectra. In several cases, a line is present in the overlap between two spectra with diffeent central wavelengths. Such lines have the same equivalent width to within the measurement errors showing the location of the line within the bandpass does not affect the continuum placement and the measurement of the equivalent width. Finally, we note (see below) that the abundances derived from weak lines of Fe I and Ti II show a scatter that is compatible with the known errors of measurement and incompatible with the presence of large systematic errors vitiating the measured equivalent widths.

The observed lines are listed in Table 2. Portions of sample spectra are shown in Figures 1, 2, and 3. In some cases, detectable lines were rejected for analysis because of blends with neighbors or the absence of high-quality spectra. Some blended lines were retained: the individual lines of the C I, O I (Fig. 2), Mg II 4481 Å (Fig. 1), and S I multiplets in the λ Boo stars are completely smeared into a single feature, but the blends are not contaminated by other elements. The total blend can be safely used to determine the elemental abundances.

Equivalent widths derived from the spectra were converted to abundances using the program WIDTH6 (R. L. Kurucz, private communication) and a model atmosphere selected from the grid used by Baschek and Slettebak (1988) and computed by them using the program ATLAS6 (Kurucz 1979). Where necessary, synthetic spectra were computed using the program MOOG (Sneden 1973) and matched to an observed spectrum. The assumption of local thermodynamic equilibrium (LTE) is adopted for the model atmospheres' construction and the analysis of the lines. Baschek and Slettebak's set of model atmospheres cover the effective temperature range ($T_{\rm eff}$) of 7500 K to 11,000 K, with all models having the surface gravity log $g = 4.0$; the λ Boötis stars and Vega are all on or near the main sequence. Models with both solar and 1/3-solar metallicities were included in the grid. Similar model atmospheres (Kurucz 1979) for surface gravities of log $g = 3.5$ and 4.0 were also used.

The defining atmospheric parameters, $T_{\rm eff}$ and g, were calcu-

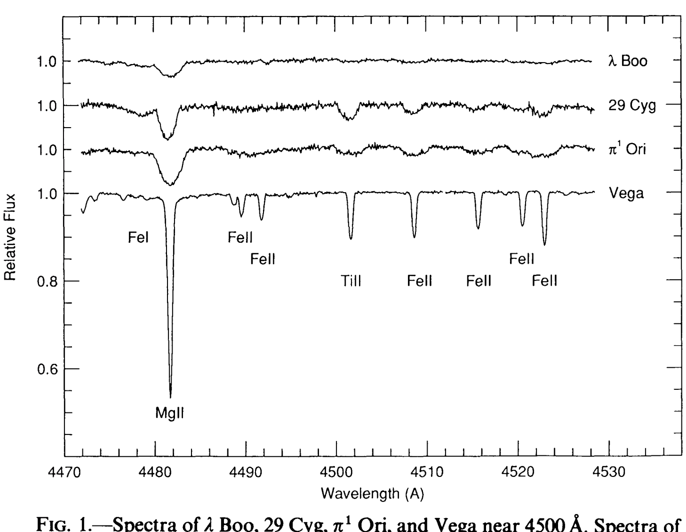

FIG. 1.—Spectra of λ Boo, 29 Cyg, π^1 Ori, and Vega near 4500 Å. Spectra of the former three stars have been displaed for clarity. The Mg II 4481 Å multiplet and assorted lines of Ti II and Fe II are identified.

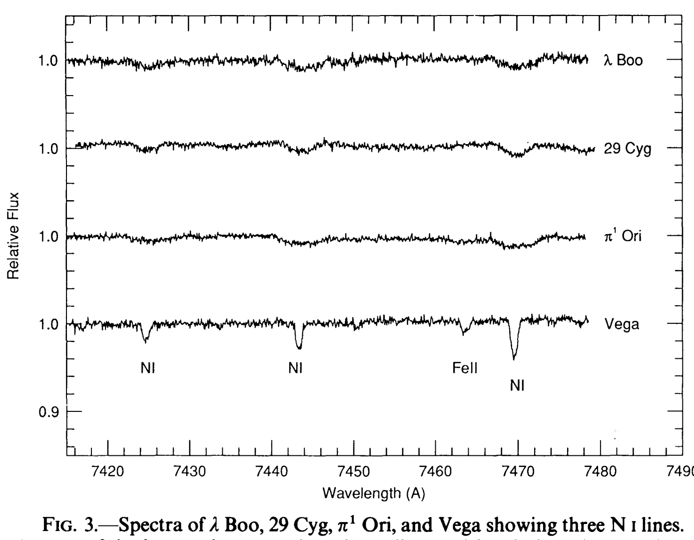

FIG. 3.—Spectra of λ Boo, 29 Cyg, π^1 Ori, and Vega showing three N I lines. Spectra of the former three stars have been displaced for clarity. The error bar (±0.25 dex) is the typical estimate (see Table 5 and text).

TABLE 2

THE OBSERVED LINES

				VEGA		λ Boo		29 Cyg		π^1 Ori	
ELEMENT (MULTIPLET)	λ (Å)	χ (eV)	log gf[a]	W_λ	log ϵ	W_λ	log ϵ	W_λ	log ϵ	W_λ	log ϵ
C I (25.02)	7108.94	8.65	−1.70	7	8.42						
C I (26)	7111.48	8.65	−1.21	12	8.38						
C I (26)	7113.18	8.65	−0.82	22	8.39	178	8.19	278	8.32	195	8.29
C I (26)	7115.19	8.65	−0.79	21	8.26						
C I (25.02)	7116.99	8.65	−0.97	27	8.54						
C I (25.02)	7119.67	8.65	−1.20	16	8.50						
N I (3)	7423.63	10.33	−0.70	18	8.03	22	8.07	18	8.01	18	7.98
N I (3)	7442.28	10.33	−0.41	27	7.97	43	8.20	24	7.88	28	7.95
N I (3)	7468.29	10.33	−0.21	39	8.01	47	8.07	31	7.85	52	8.15
O I (10)	6155.99	10.74	−0.66								
O I (10)	6156.79	10.74	−0.44	122	8.74	90	8.48	78	8.47	90	8.48
O I (10)	6158.19	10.74	−0.29								
Na I (1)	5889.95	0.00	0.11	112	6.32	78	4.97	92	4.62	140	5.79
Na I (1)	5895.92	0.00	−0.19	98	6.42	63	5.10	67	4.75	118	5.78
Mg II (1)	9217.40	8.65	0.26	99	6.59			36	5.77		
Mg II (1)	9243.40	8.65	−0.04	80	6.65						
Mg II (4)	4481.13	8.86	0.75	296	6.89	89	5.58	126	5.92	192	6.19
Mg II (4)	4481.33	8.86	0.57								
S I (1)	9212.91	6.52	−0.42	120	7.42			200	7.23		
S I (1)	9237.49	6.52	−2.00	74	7.30			163	7.17		
S I (8)	6757.16	7.87	−0.29					15	6.73	13	6.87
S I (3)	10459.46	6.86	−1.52	40	7.10						
S I (3)	10455.47	6.86	0.26	73	7.00			330:	6.8 :	430	7.55
S I (3)	10456.79	6.86	−0.44								
Ca I (2)	4226.73	1.90	0.24	60	5.85						
Ca I (3)	6122.22	1.89	−0.32	5	6.04	<8	<5.29	23	5.34	<11	<5.47
Ca I (3)	6162.17	1.90	−0.09	5	5.95	<8	<5.04	23	5.13	<12	<5.29
Ca II (1)	3933.68	0.00	0.13	656	5.84	379	4.39	1135	5.07	672	5.01
Ti II (11)	4012.37	0.57	−1.61	28	4.45						
Ti II (41)	4290.22	1.16	−1.12	63	4.86						
Ti II (19)	4443.80	1.08	−0.70	63	4.37	11	2.80	70	3.62		
Ti II (31)	4468.49	1.13	−0.60	67	4.36	19	3.02	73	3.83	55	3.60
Ti II (31)	4501.27	1.12	−0.75	56	4.35	9	2.77	49	3.42	39	3.58
Ti II (50)	4533.97	1.24	−0.77	80	4.77	16	3.18	73	3.83	63	3.99
Ti II (50)	4563.76	1.22	−0.96	55	4.64			45	3.65	42	3.90
Ti II (82)	4571.97	1.57	−0.53	74	4.55	12	3.01	65	3.74	51	3.83
Fe I (4)	3920.26	0.12	−1.75	22	7.10			39	5.89		
Fe I (4)	3922.91	0.05	−1.65	22	6.96			40	5.76		
Fe I (4)	3930.30	0.09	−1.59	32	7.13						
Fe I (278)	3997.39	2.73	−0.39	12	7.01					20	6.50
Fe I (43)	4005.25	1.56	−0.61	38	7.14	17	5.78	47	5.90	40	6.32
Fe I (43)	4045.82	1.48	0.28	70	6.69	55	5.51	112	5.84	101	6.19
Fe I (43)	4063.60	1.56	0.07	61	6.81	38	5.54	94	5.81	88	6.27
Fe I (43)	4071.74	1.61	−0.02	53	6.81	36	5.63	89	5.88	73	6.20
Fe I (42)	4202.03	1.48	−0.71	26	6.96	9	5.50	46	5.90	36	6.30
Fe I (152)	4235.94	2.43	−0.34	16	6.90			30	5.95		
Fe I (152)	4260.48	2.40	0.02	33	6.96	17	5.74	47	5.87	54	6.47
Fe I (71)	4282.41	2.18	−0.81	14	7.15						
Fe I (350)	4466.55	2.83	−0.59	9	7.12						
Fe II (27)	4233.17	2.58	−2.00	90	6.91	36	5.64	74	5.98	99	6.49
Fe II (37)	4472.92	2.84	−3.43	8	6.88						
Fe II (38)	4508.28	2.86	−2.21	51	6.48	12	5.47	31	5.76	46	6.19
Fe II (37)	4515.34	2.84	−2.48	42	6.80			19	5.77	32	6.25
Fe II (37)	4520.23	2.81	−2.60	41	6.88						
Fe II (38)	4522.63	2.84	−2.03	64	6.69			34	5.62		
Fe II (38)	4541.52	2.86	−3.05	23	7.02			10	6.04	19	6.53
Fe II (38)	4576.33	2.84	−3.04	24	7.02			8	5.89	20	6.54
Fe II (37)	4582.84	2.84	−3.10	14	6.82						
Fe II (38)	4583.83	2.81	−2.02	88	7.00			58	5.93		
Fe II (37)	4629.34	2.81	−2.37	51	6.81	17	5.75	32	5.89		
Sr II (1)	4077.71	0.00	0.15	47	1.98	39	0.86	123	1.60	56	1.16
Sr II (1)	4215.52	0.00	−0.16	35	2.07	18	0.74	103	1.54	46	1.32

[a] Sources of gf-values: C I: See text; N I: See text; O I: Lambert, Roby, and Bell 1982; Na I. Mg II, Ca I, Ca II, Sr II: Wiese and Martin 1980; S I: Lambert and Luck 1978; Ti II, Fe I, Fe II: Fuhr, Martin, and Wiese 1988.

TABLE 3

EFFECTIVE TEMPERATURES AND SURFACE GRAVITIES

	π^1 Ori		λ Boo		29 Cyg	
REFERENCE	$T_{\rm eff}$ (K)	log g (cm s^{-1})	$T_{\rm eff}$ (K)	log g (cm s^{-1})	$T_{\rm eff}$ (K)	log g (cm s^{-1})
Relyea and Kurucz 1978	8750	4.0	8650	4.0	8200	4.0
Moon and Dworetsky 1985	8800	4.2	8700	4.2	8000	3.9
Lester, Gray, and Kurucz 1986	8725	4.0	8600	4.0	8000	3.7
Baschek and Searle 1969	8550	4.0	8400	4.0	8000	3.9
Oke 1967	8250	4.0	8250	3.9	7750	3.8
Baschek and Slettebak 1988	8750	4.0	8800	4.0	8150	4.0
Adopted	8750	4.0	8650	4.0	8100	4.0

lated for each star from their Strömgren photometric indices (Hauck and Mermilliod 1980), which are listed in Table 1. The indices were dereddened as outlined by Crawford (1970). The atmospheric parameters (Table 3) were then determined in three ways from the Strömgren indices. Relyea and Kurucz (1978) calculated theoretical Strömgren colors using the Kurucz (1979) grid of model atmospheres. These were calibrated to a standard system using a model atmosphere for Vega of $T_{\rm eff} = 9400$ K, log $g = 3.95$. Predictions for models at 1/10 and 1/100 solar abundances are included. In determining $T_{\rm eff}$ and log g for our stars from their $(c_1)_0$ versus $(b-y)_0$ graph, we used the solar abundance models, but considered the metal-deficient models in estimating the errors.

Moon and Dworetsky (1985) found that adjustments are necessary to the Relyea and Kurucz synthetic colors to bring them into agreement with observed indices. They also apply a scaling correction to the synthetic β indices calculated by Schmidt (1979) from Kurucz models. The adjustments are derived empirically by comparing the theoretical and observed color indices of eclipsing and visual binaries of known masses and radii. The results are presented in (β, c_0) or (a_0, r^*) grids to determine $T_{\rm eff}$ and log g (the indices are defined in their paper). The roles of the β and c_0 indices, one as a luminosity indicator (and hence, surface gravity) and the other as a temperture parameter, become ambiguous between spectral types B9–A3. They find the a_0 and r^* parameters can overcome this problem. However, the synthetic β index appers insensitive to the bulk metallicity of a model atmosphere which makes it an ideal index to use for the λ Boo stars. We use the (a_0, r^*) grid primarily in estimating the $T_{\rm eff}$ and log g, but 29 Cyg's parameters were found from the (β, c_0) for $T_{\rm eff} \leq 8500$ K.

Lester, Gray, and Kurucz (1986) have recalibrated theoretical photometric indices derived from Kurucz atmospheres using the ultraviolet and visual energy distributions from five spectrophotometric standard stars (γ Gem, α CMi, β Leo, η UMa, and Vega). Although there are systematic differences between their results and those of Relyea and Kurucz, the newer results for $T_{\rm eff}$ and log g are not significantly different from the other predictions.

The three estimates of $T_{\rm eff}$ and log g for the λ Boo stars were averaged to obtain our best estimates. These are compared in Table 3 to those of Baschek and Slettebak (1988). Their $T_{\rm eff}$'s were estimated from $T_{\rm eff}$ versus $(b-y)$ and $(b-y)$ versus $(B-V)$ relations with $T_{\rm eff}$'s taken from the calibrations by Code *et al.* (1976) and Böhm-Vitense (1981). We also give in Table 3 the $T_{\rm eff}$ and log g adopted by Baschek and Searle (1969) for their abundance analysis. They and Oke (1967—see Table 3) derived these estimates from the continuum energy distribution across the optical region as well as Balmer line profiles; observations were fitted to predictions based on model atmospheres computed by Mihalas (1965, 1966). A detailed study of Vega was made by Dreiling and Bell (1980), and for this study, their values for the atmospheric parameters have been adopted: $T_{\rm eff} = 9650$ K and log $g = 3.95$, although our model corresponded to log $g = 4.0$.

The vast majority of the measured lines are insensitive to the adopted microturbulent velocity. A value of 3 km s^{-1} was adopted for the λ Boo stars. For Vega, we adopted the value of 2.0 km s^{-1} found by Lambert, Roby, and Bell (1982). For the strongest lines, radiative and Stark broadening was taken into account using the values provided by the program WIDTH6, except for Ca II and Mg II to be discussed later.

The adopted *gf*-values are listed in Table 2 with the measured equivalent widths and the abundances derived using the adopted model atmospheres. In general, the adopted *gf*-value is taken from a recent critical compilation of theoretical and experimental estimates—see references listed in Table 2. For C I, our values are based on the "solar" *gf*-values given by Lambert, Roby, and Bell (1982) which assumed a solar carbon abundance of log ϵ(C) = 8.67 (Lambert 1978). A redetermination of this abundance has lowered it to log ϵ(C) = 8.56 (see Anders and Grevesse 1989), and hence, the solar *gf*-values are here increased by 0.11 dex over the values given by Lambert, Roby, and Bell (1982).

II. THE CHEMICAL COMPOSITIONS

The chemical compositions are summarized in Table 4. We comment first on particular aspects of the line selection for some individual elements, give a discussion of random and systematic errors of measurement, and conclude with a comparison between our results and those previously published from optical and ultraviolet spectra.

The carbon abundance is provided by the multiplet near 7115 Å. Individual lines are well resolved in the spectrum of the slowly rotating Vega ($v \sin i = 16$ km s^{-1}) and give a consistent set of abundances for which the mean ([C] = −0.14, relative to the Sun) is in good agreement with that found independently by Roby and Lambert (1990; [C] = −0.11). Individual lines are not resolved in the spectra of the λ Boo stars.

Nitrogen abundances are estimated from the triplet near 7450 Å. When differences in the adopted *gf*-values are taken into account, our N abundance for Vega is within 0.03 dex of the value given by Roby and Lambert (1990). Spectra near 8690 Å were obtained to provide additional N I lines for the λ Boo stars, but these lines were badly blended with the wings of much stronger Paschen lines. A careful synthesis of this region may provide an estimate of the N abundance.

The 6158 Å O I multiplet is rotationally blended in the λ Boo

TABLE 4

ELEMENTAL ABUNDANCES

Species	At. No.	FIP (eV)	SIP (eV)	λ Boo log ϵ	λ Boo [ϵ]	29 Cyg log ϵ	29 Cyg [ϵ]	π^1 Ori log ϵ	π^1 Ori [ϵ]	Vega log ϵ	Vega [ϵ]	Sun[a] log ϵ
C I	6	11.26	24.38	8.19	−0.37	8.32	−0.24	8.29	−0.27	8.42	−0.14	8.56
N I	7	14.53	29.60	8.11	+0.06	7.91	−0.14	8.03	−0.02	8.00	−0.05	8.05
O I	8	13.62	35.12	8.48	−0.45	8.47	−0.46	8.48	−0.45	8.74	−0.19	8.93
Na I	11	5.14	47.29	5.03	−1.30	4.68	−1.65	5.79	−0.54	6.37	+0.04	6.33
Mg II	12	7.65	15.04	5.58	−2.00	5.85	−1.73	6.19	−1.39	6.71	−0.87	7.58
S I	16	10.36	23.33	...	...	6.98	−0.23	7.21	0.00	7.21	0.00	7.21
Ca I	20	6.11	11.87	<5.17	< −1.19	5.24	−1.12	<5.38	< −0.98	5.95	−0.41	6.36
Ca II	20	...	...	4.39	−1.97	5.07	−1.29	5.01	−1.35	5.84	−0.52	6.36
Ti II	22	6.82	13.58	2.96	−2.03	3.64	−1.35	3.78	−1.21	4.55	−0.44	4.99
Fe I	26	7.87	16.16	5.62	−2.05	5.87	−1.80	6.32	−1.35	6.98	−0.69	7.67
Fe II	26	...	...	5.62	−2.05	5.86	−1.81	6.40	−1.27	6.85	−0.82	7.67
Sr II	38	5.70	11.03	0.80	−2.10	1.57	−1.33	1.24	−1.66	2.03	−0.87	2.9

[a] Anders and Grevesse 1989.

stars. Our abundance for Vega is smaller by 0.10 dex than the value given by Roby and Lambert (1990). The origin of this difference is in the measured equivalent widths.

Sodium abundances are derived from the NaD lines. Telluric H_2O lines were removed from the spectra using the spectrum of a very rapidly rotating hot star. Although the Na deficiency of the λ Boo stars is clearly established, the (LTE) abundances are moderately sensitive to the adopted microturbulent velocity.

The Mg abundance is determined from the Mg II 4481 Å feature whose two components are unresolved even in the slow rotator Vega. In the spectra of the rapidly rotating λ Boo stars, the Mg II feature is slightly blended with weak Fe I lines whose contribution to the equivalent width was subtracted to obtain the Mg II feature's equivalent width. Our final Mg abundances from synthetic spectra of the 4481 Å feature confirm that Mg is underabundant in λ Boo stars, as was suggested by Morgan, Keenan, and Kellman's (1943) discovery of a weak Mg II feature. Our abundance for Vega is within 0.11 dex of that given by Gigas (1988) from his LTE analysis for this feature. For Vega and 29 Cyg, the Mg abundances are also calculated from measurements of weaker Mg II lines near 9220 Å which result in slightly lower abundances.

As the discussion in § IV will reveal, an effort to determine the S abundances in the λ Boo stars was started after the abundance analyses were otherwise complete. Takada-Hidai (1990) has recently determined Vega's S abundance from the triplet of red S I lines near 6750 Å. On our CCD spectra of 29 Cyg and π^1 Ori, these lines are very weak such that only the line at 6757 Å is detectable above the noise level. CCD spectra of stronger S I lines near 9220 and 10456 Å were obtained for 29 Cyg and π^1 Ori, as well as Vega. A considerable number of telluric H_2O lines contaminate the former region. While these may be removed by ratioing the spectra to those of hot rapidly rotating stars observed at similar air masses, the resultant spectra are not the intrinsic spectra because a broad Paschen line falls in this region. Equivalent widths of the S I (and Mg II) lines were measured relative to the local continuum which is defined by the ratioing of the two Paschen lines. Two of the three S I lines are well-defined, but the third line is too close to the core of the Paschen line to be measurable. The 10456 Å triplet is seen in the spectrum of π^1 Ori as a single blended feature. The blended S I triplet is present in 29 Cyg, but a measurement of its equivalent width is uncertain due to imperfect cancellation of the large amplitude of interference fringes introduced by the CCD. Our conclusion that S has a near-normal abundance in the λ Boo stars is in mild disagreement with Baschek and Slettebak's (1988) measurements of a single ultraviolet S I line that suggest S is slightly underabundant.

The Ca abundances are derived primarily from the CA II K line, although two Ca I lines were measured in Vega and 29 Cyg. The K line is between the Balmer Hϵ and H8 lines whose contribution to the opacity is included in WIDTH6. Our Ca abundances are similar to those obtained by Baschek and Searle (1969), who also used the K line in their abundance determination: they found [Ca/H] = (−1.4, −0.9, −1.0) and we find (−1.4, −0.8, −0.8) relative to Vega for the trio (λ Boo, 29 Cyg, π^1 Ori). The Ca I resonance line at 4226 Å, which is observable in Vega, is blended with Fe I lines in the λ Boo stars. The Ca I lines at 6122 Å and 6162 Å were measured in Vega and 29 Cyg but are not detectable in λ Boo and π^1 Ori. The abundances (or the upper limits) given by the Ca I lines are consistent with those derived from the K line.

Several Ti II lines were measured for each λ Boo star. The line-to-line spread (±0.2 dex) of the Ti abundances is consistent with the measurement errors of these shallow lines and the uncertainties of the *gf*-values. The latter are removed as a source of error when Ti abundances relative to Vega are considered; the non-LTE effects are also partially cancelled when these relative abundances are computed. Then the line-to-line spread is reduced to ±0.12 dex for λ Boo and only ±0.03 dex for π^1 Ori, for which the Ti II lines are measureable with greater accuracy than for λ Boo. Our Ti abundance for Vega is within 0.05 dex of that obtained by Gigas and quoted by Lemke (1989—see also earlier determinations by Dreiling and Bell 1980; Sadakane and Nishimura 1981). The Ti abundances of the λ Boo stars are substantially smaller than those given by Baschek and Searle (1969), i.e., [Ti/H] relative to Vega is (−1.6, −0.9, −0.8) from our analyses, but (+0.2, −0.5, −0.5) from Baschek and Searle, for the trio (λ Boo, 29 Cyg, π^1 Ori). These discrepancies arise almost entirely from differences in the measured equivalent widths. Lambda Boo is the most outstanding case. Baschek and Searle reported a measurement for one line at 4443 Å and upper limits for six additional lines. Their measurement of the 4443.3 Å line (W_λ = 110 mÅ) is an order of magnitude larger than our measurement (W_λ = 11 mÅ). The extreme weakness of Ti II lines in the spectrum of λ Boo is evident from Figure 1. The smaller differences between our and their [Ti/H] estimates for 29 Cyg and π^1 Ori reflects the closer agreement for the equivalent widths; e.g., the equivalent widths of three lines measured by us for π^1 Ori are only about 60% smaller than those listed by Baschek and Searle.

Our spectra provide a good selection of Fe I and Fe II lines. The line-to-line scatter in the abundances (log ϵ) is about ± 0.15 dex, a value that is plausibly attributed to the combination of errors associated with the *gf*-values and the equivalent widths. Unlike the example provided by the Ti II lines, the ranges are not decreased when abundances relative to Vega are considered. This difference is not surprising because the adopted *gf*-values for Fe I and, perhaps, Fe II are often of greater accuracy than those for Ti II. Our abundances for Vega are very similar to the LTE abundances given recently by Gigas (1986) from an independent selection of the *gf*-values and lines: log ϵ(Fe) = 6.98 (Fe I) and 6.85 (Fe II) from our study and log ϵ(Fe) = 6.92 (Fe I) and 7.01 (Fe II) from Gigas (1986). Similar Fe deficiencies for Vega have been reported by other authors (Dreiling and Bell 1980; Sadakane and Nishimura 1981; Lane and Lester 1984). A thorough non–LTE study of Fe I and Fe II lines in Vega was reported by Gigas (1986): the non–LTE abundance for his selections of lines is raised by 0.3 dex for Fe I and reduced by a mere 0.02 dex for Fe II lines. Calculations reported by Lemke (1989) show that the non–LTE effects are insensitive to effective temperature and so should be similar for the λ Boo stars. Then, our abundances after correction are mildly discordant for Fe I and Fe II: log ϵ (Fe) $\simeq$ 7.3 and 6.9 for Fe I and Fe II, respectively. The non–LTE calculations for Vega (Gigas 1986) give a similar disagreement (7.2 vs. 7.0). For Lemke's (1989) sample of 16 B9.5 V to A2 V stars, the mean difference for the non–LTE abundances is 0.16 dex in the above sense. For our present purpose, these small differences may be ignored. Systematic errors in the $T_{\rm eff}$ scale, the *gf*-values, and the non–LTE calculations seem to be likely contributors to the differences.

In contrast to the good agreement between our and other determinations of Vega's Fe abundance, our results for the λ Boo stars show them to be more metal-poor than suggested previously. For example, Baschek and Searle (1969) reported [Fe/H] from Fe II lines to be (-0.5, $+0.1$ and -0.5) relative to Vega, but we obtain (-0.6, -1.3, and -1.1) for π^1 Ori, λ Boo, and 29 Cyg, respectively. As was pointed out in the discussion of the Ti II lines, the principal reason for our lower abundances is that the equivalent widths measured by us are smaller; for example, four Fe II lines were reported by Baschek and Searle to have equivalent widths of 80 to 100 mÅ in 29 Cyg, but our spectra yield measurements of 19 to 58 mÅ. For λ Boo, our measured equivalent widths for Fe I and Fe II lines are all smaller than the detection limit ($W_\lambda \sim 80$ mÅ) set by Baschek and Searle's photographic spectra. It is quite apparent from our spectra that the λ Boo stars span a range in metallicity.

Strontium was included in our survey as a monitor of the *s*-process abundances that might provide clues to the origin of the λ Boo stars. The Sr II resonance lines at 4077 and 4215 Å were measured in all stars. The equivalent width of a Sr II line is quite similar in the three λ Boo stars and Vega, but the Sr abundance is considerably higher in the latter star because its effective temperature is about 1000 K hotter. A search was conducted for the 4554 and 4525 Å lines of Ba II; Ba is a much heavier *s*-process element than Sr. The latter line is too weak for reliable measurement in Vega. The former, which has an equivalent width of only 11 mÅ in Vega, yields an abundance log ϵ (Ba) = 1.5 that is consistent with an independent (LTE) determination [log ϵ (Ba) = 1.6 $\pm$ 0.2, Gigas 1988]. Unfortunately, the 4554 Å line in the λ Boo stars is blended with stronger Cr II and Fe II lines and is therefore ill-suited to an abundance analysis.

Readily estimated measurement errors are summarized in Table 5 for π^1 Ori. These errors refer to the defining parameters of the model atmosphere ($T_{\rm eff}$, g, ξ, and the assumed metallicity Z) and the equivalent widths. Similar estimates were found for λ Boo and 29 Cyg; the sensitivity to the adopted microturbulence is generally greater for π^1 Ori because the observed lines are stronger, e.g., $\Delta \log \epsilon = -0.08$ for Na I in λ Boo. The sensitivity of the S abundance to the adopted microturbulence is derived for the strong near-infrared lines; the red lines are insensitive to the microturbulence. Vega is a hotter star with many stronger lines, and the abundances of Na and Mg are significantly more sensitive to the adopted microturbulence than suggested by Table 5. Sr is also more sensitive to microturbulence in the cooler star 29 Cyg, but other estimated errors are similar to those in the table. Analysis of the Mg II and Ca II lines is sensitive to the adopted damping constants. WIDTH6 provides estimates of the radiative and Stark contributions to the damping constant. We made independent estimates using the best available transition probabilities for the radiative contribution and experimental line widths (Konjevic and Wiese 1979) for the Stark width. For Ca II, the Ca abundance was reduced by about 0.10 dex when our estimates of the damping constant were adopted. For Mg II, the reduction was smaller (0.02 to 0.07 dex). Errors in the adopted *gf*-values must be included in estimating the total overall uncertainty in an abundance ϵ. For several elements (Na, Mg, Ca, Sr), the selected lines have well-determined *gf*-values based on accurate measurements of radiative lifetimes; the uncertainties are less than ± 0.10 dex and a negligible contributor to the total uncertainty. The C I *gf*-values are "solar" values, and hence, tied to the adopted solar C abundance and the model solar atmosphere. As long as the stellar C abundances are interpreted relative to the solar value, the C I *gf*-values probably contribute between ± 0.05 and ± 0.10 dex uncertainty to the stellar abundances. The N I *gf*-values are the "best" values from Grevesse *et al.* (1990) who give a compilation of estimates including results of several recent theoretical calculations. These "best" values for the $3s$–$3p$ transition array to which our selected lines belong yield a solar N abun-

TABLE 5

ESTIMATED ABUNDANCE UNCERTAINTIES ($\Delta \log \epsilon$) FOR π^1 ORIONIS

Source	C I	N I	O I	Na I	Mg II	S I	Ca I	Ca II	Ti II	Fe I	Fe II	Sr II
$T_{\rm eff}$ (+200 K)	0.06	−0.01	−0.01	0.15	−0.02	0.06	0.18	0.10	0.06	0.14	0.06	0.17
ξ (+1.0 km s^{-1})	−0.02	−0.03	−0.03	−0.20	−0.12	−0.42	−0.01	−0.03	−0.06	−0.13	−0.08	−0.08
Z (−0.5 dex)	−0.04	−0.00	0.00	−0.03	0.02	0.05	−0.06	−0.04	−0.03	−0.06	−0.03	−0.05
Line measures (±)	0.05	0.10	0.05	0.07	0.02	0.10	...[a]	0.02	0.10	0.10	0.10	0.10
log g (+0.5 dex)	0.03	−0.07	−0.07	0.08	−0.07	0.07	0.15	0.03	−0.10	0.10	−0.11	−0.01

[a] The uncertainty for Vega and 29 Cyg (where the lines are measurable) is about ±0.15.

dance [log ϵ (N) = 8.00 ± 0.10]. The recommended solar abundance, log ϵ (N) = 8.03, is based on these and other N I lines and the infrared vibration-rotation lines of the NH molecules. We estimate that the stellar abundances are subject to a maximum uncertainty of ±0.10 dex contributed by the N I *gf*-values. A similar uncertainty applies to the O abundance as estimated from the theoretical and experimental O I *gf*-values and the solar 6158 Å lines interpreted using the accurate O abundance derived from the pure rotation and the vibration-rotation OH lines and the [O I] lines. For the remaining species (e.g., Fe I, Fe II, and Ti II), the *gf*-related abundance uncertainties are probably in the range ±0.05 to ±0.20. Of course, the *gf*-value is of no concern when the stellar relative abundance is estimated from measurements of the same line in the λ Boo star and Vega or the Sun. Relative abundances computed from samples comprising different selections of lines are subject to a small uncertainty from the *gf*-values.

Although our analysis generally confirms earlier reports that the λ Boo stars are metal poor (Na–Sr) with approximately normal abundances of C, N, and O, there are several novel results that deserve some comment. It is clear from our study that the λ Boo stars span a range in metal-abundance, but all have similar and nearly normal abundances of C, N, O, and S. Of the present trio, λ Boo is the most metal-poor ([Fe/H] = −2.0) and π^1 Ori is the least metal-poor ([Fe/H] = −1.3). These relative abundances, are given with respect to the Sun, not the "standard" star Vega for which [Fe/H] = −0.7. The Fe abundances and those of other elements from Na to Sr differ significantly from those published by Baschek and Searle (1969) in their analysis of photographic optical spectra and by Baschek and Slettebak (1988) from an analysis of *IUE* high-resolution spectra. In the former study, the Fe abundances relative to Vega were given as (+0.1, −0.5, −0.5) for the trio (λ Boo, 29 Cyg, π^1 Ori), but we obtain (−1.3, −1.1, −0.6), i.e., λ Boo and 29 Cyg are both much more metal-poor than suggested by Baschek and Searle (1969). As noted earlier in our discussion of Ti and Fe, these large differences are traceable to the equivalent widths. The large equivalent widths listed by Baschek and Searle are inconsistent with our measurements from spectra of superior quality.

Baschek and Slettebak (1988) discuss the *IUE* spectra by presenting diagrams showing observed and predicted equivalent widths of particular lines or blends as a function of the $(b-y)$ color (i.e., the effective temperature). The stars investigated by them include our three λ Boo stars and Vega, as well as other established or suspected λ Boo and standard stars. The predicted (LTE) equivalent widths were obtained using the grid of model atmospheres from which our models were selected. Predictions were made for three or four abundances of the element responsible for the measured feature. By inspection of the diagrams for the individual features, the abundances may be obtained. Five Fe II features give [Fe/H] relative to Vega as (−0.5, −0.5, −0.2) for (λ Boo, 29 Cyg, π^1 Ori) and these abundances are systematically higher than our values of (−1.3, −1.1, −0.6). Baschek and Slettebak's abundances are similarly systematically higher for Mg, the other metal which is common to both studies. The two studies are in slightly better agreement when the relative abundance is given with respect to standard stars of the same temperture as the λ Boo stars. Absolute abundances derived by comparing the observed and predicted equivalent widths show a large scatter with mean values disagreeing sharply with our estimates; e.g., the five Fe II ultraviolet features measured in λ Boo give log ϵ (Fe) of 8.0, 7.2, 6.5, 7.0, and 7.2 for a mean value of 7.2 in contrast to our result of 5.6. Since the lines measured from *IUE* spectra are strong, the derived abundance is sensitive to the adopted atmospheric structure, the damping constants, non–LTE effects, and to blends. Quite possibly, the non–LTE effects are far from cancelled in differential analyses when, as in the case, a line has different equivalent widths in the λ Boo and standard stars.

The first hint that the light elements (C, N, and O) do not share the metals' large underabundances in λ Boo stars was offered by Kodaira (1967; see also Baschek and Searle's remarks in Bowen 1963) who considered the strong O I 7770 Å triplet in λ Boo and π^1 Ori. Baschek and Searle's (1969) observations and analysis of this triplet gave [O/H] = +0.6, +0.2, and −0.2 (relative to Vega for λ Boo, 29 Cyg, and π^1 Ori, respectively). Observations of the strong C I 9100 Å lines led Lambert, McKinley, and Roby (1986) to conclude that carbon was not underabundant in λ Boo stars. Baschek and Slettebak's (1988) analysis of *IUE* spectra led them to give the following mean abundances for their sample of λ Boo stars relative to standard stars: [C/H] = +0.3 from one C I and one C II feature, [N/H] = +0.2 from four N I lines, and [O/H] = +0.1 from one O I line. When Baschek and Slettebak's diagrams are inspected, one finds [C/H] (relative to Vega) for our three λ Boo stars to be slightly less than the above average, and [N/H] = +0.4 for all three stars. Little can be said concerning oxygen because the O I 1302 Å line was not measured in λ Boo and 29 Cyg, and the measurement for π^1 Ori is denoted as uncertain. Our results are in fair agreement with these estimates. As noted earlier, the *IUE* lines are very saturated and sensitive to the non–LTE effects. It should also be noted that *IUE* spectra were not measured for standard stars at the same $(b-y)$ color as our λ Boo stars. Since the chosen standard stars were hotter than the λ Boo stars, systematic errors may vitiate the differential abundance analysis. Nonetheless, Baschek and Slettebak provided for the first time convincing evidence that all three of the common light elements (C, N, and O) have essentially normal abundances in contrast to the underabundances of the investigated metals. Our analysis confirms this result. Corrections for non–LTE effects for the light elements will probably not change this result significantly (see Lambert, Roby, and Bell 1982 for a brief discussion).

IV. POSSIBLE ORIGINS OF λ BOOTIS STARS

After the pioneering quantitative spectroscopy of λ Boo stars in the 1960s, the enigma posed by their atmospheric compositions seems to have been forgotten or dismissed as either uninteresting or too challenging to crack with the available data on compositions and other properties. Upon completion of the ultraviolet studies (Bashek and Slettebak 1988) and our analysis, it seems timely to search afresh for the origin of these metal-poor A stars.

Pertinent information to consider in the search is the stars' membership among Population I stars. This membership is soundly based on reliable trigonometric parallaxes that place the stars on or near the normal main sequence, radial velocities, and proper motions that show the stars to have the kinematics expected of young disk stars, and colors and hydrogen line profiles expected of young main-sequence stars rather than the more luminous, less massive A stars of Population II. Clear assignment to Population I dissuaded us from entertaining speculations spawned by a naive identification of the stars as Population II (i.e., metal-poor) stars that have raised their C, N, and O abundances through severe mass loss and mixing of C and O from He-burning layers and of N from H-burning layers. Furthermore, as noted by Baschek and Slettebak (1988), the similar abundances (relative to solar) of C, N, and O in λ Boo stars places unrealistic constraints on the relative proportions of the H and He-burning products.

The hypothesis of element segregation through diffusion (Michaud 1970, 1987) is certainly the leading contender to explain the various families of chemically peculiar stars on the upper main sequence. An attempt was made by Michaud and Charland (1986) to explain the λ Boo stars as the result of diffusion in the presence of mass loss. An encouraging measure of agreement between the predicted and observed compositions was claimed, but we must note that major disagreements exist between the predictions and the more complete and reliable compositions now available: (1) the diffusion calculations account for the modest (a factor of 3) metal-deficiency suggested by the early abundance analyses, but not the much larger deficiencies found in this study, e.g., [Fe/H] ~ -1.3 to -2.0; (2) the calculations show that large variations in abundance are expected from metal to metal in contrast to the rather uniform deficiencies reported here; and (3) the available calculations for C and O indicate that the light elements may experience similar deficiencies to the metals, but the recent analyses show that normal C, N, and O abundances are a firm characteristic of λ Boo stars. Our conclusion is that the published study of diffusion does not account for the λ Boo stars. Two other points may be mentioned. First, the composition of λ Boo stars is quite unlike that of any one of the several well-observed families of chemically peculiar stars. It seems, therefore, that if diffusion is to account for the λ Boo stars as well as the latter families, one must identify a characteristic such as rotation or mass loss that, in association with diffusion, leads to the distinctive composition of a λ Boo star. Then a complete explanation based on diffusion should explain why this small group of stars experience mass loss. Second, as Baschek and Slettebak (1988) remark, meridional flows induced in rapidly rotating λ Boo stars may inhibit the establishment of abundance anomalies. In view of these various difficulties, we consider that diffusion is unproven as the key factor behind the λ Boo stars.

Clues to the origins of peculiar stars may be obtained by finding stars with similar compositions. Our search led quickly to the realization that the λ Boo stars have compositions similar to those reported recently for a class of luminous halo stars called post–AGB stars that are widely identified as Population II stars on account of their low metal abundance and high galactic latitude. These stars are supposed to be evolving from the AGB to become the hot central star of a planetary nebula created from the circumstellar envelope known to surround the present post–AGB star. The lack of obvious overabundances of the *s*-process elements in the post–AGB stars suggests that if the progenitors evolved from the AGB, they left it before the onset of the thermal pulses that are expected to enhance the atmosphere in carbon and the *s*-process elements. The similar compositions of λ Boo and post–AGB stars was noted first by Lambert, Hinkle, and Luck (1988) in their analysis of HR 4049, a high Galactic latitude metal-poor A-type supergiant. (Parthasarathy, Pottasch, and Wamsteker [1988] noted that the ultraviolet spectra of post–AGB and λ Boo stars have a broad unidentified absorption feature in common: a feature at 1657 Å with a width of about 100 Å.) The post–AGB and λ Boo stars differ substantially in luminosity.

A comparison of the compositions of our trio of λ Boo stars, Vega, and the post–AGB star, HD 46703 (Luck and Bond 1984; Bond and Luck 1987), is provided by Figures 4 and 5. The similarities in composition are striking. Baschek and Slettebak (1988) who were the first to remark upon the similar pattern of the abundance anomalies for Vega and the λ Boo stars suggested that the former could be considered to be a "mild, non-rotating λ Boo star." Sulphur abundances would seem to be the key to an explanation for these peculiar compositions. Our analysis shows that S has a nearly normal abundance in the λ Boo stars. A recent determination of the sulpfur abundance in Vega and HR 4049 (Takada-Hidai 1990) also shows that the S abundance is close to solar. While one can concoct recipes involving ad hoc mixtures of H-buring and He-burning products to account for the normal abundances of C, N, and O, nucleosynthesis of S necessarily leads to synthesis of other elements with a similar atomic number. In all of the examined stars, S stands apart from neighboring elements and suggests that the peculiar compositions are not primarily a result of nucleosynthesis and mixing. The similar chemical compositions may indicate either that the post–AGB stars are evolved or descended from the λ Boo stars, or a similar physical effect such as diffusion has operated in the two unrelated groups of stars.

The post–AGB stars have outstanding infrared excesses attributed to circumstellar dust grains (see, for example, Parthasarathy and Pottasch 1986). Vega also has an infrared excess from circumstellar dust (Aumann *et al.* 1984). Sadakane and Nishida (1986) analysed *IRAS* observations to show that λ Boo and π^1 Ori have large excesses at 60 μm. These results suggest that there may be a link between the presence of circumstellar dust and the unusual composition of the atmospheres. Suspicions of a link are further stimulated when it is realized that the elements so underabundant in these stars are just those depleted by large factors in the interstellar gas. The underabundances of elements in the interstellar gas are attributable to their incorporation into the interstellar grains and ices on the surfaces of the grains. Our hypothesis is that a dominant component of the atmosphere of a λ Boo star is gas accreted from a circumstellar nebula without substantial accretion of the grains. We assume that gas in a dusty circumstellar nebula will show similar depletions to those observed for interstellar gas. Cardelli's (1984) analysis of gas in the circumstellar nebula around α Sco shows a pattern of depletions similar to the interstellar pattern. Alternatively, grains may have condensed in and been expelled from the atmosphere leaving the gas which is now present in the atmosphere depleted in those elements that make up the grains.

It is a characteristic of the interstellar gas that different metals are depleted by different amounts; see, for example, Morton (1974) for the composition of the clouds in front of ζ Oph where Ca and Fe are underabundant by factors of $10^{-3.7}$ and $10^{-1.7}$, respectively. At first sight, such an element-to-element variation seems at odds with the rather uniform deficiencies reported here for the λ Boo stars. However, the simplest possible model reconciles these two distinctly different patterns. We suppose that the observed atmosphere consists of a mix of interstellar gas and normal (i.e., solar) gas in the proportions f to $1-f$; mixing following accretion may be driven by the meridional currents in these rapidly rotating stars. Then, an elemental abundance $\epsilon(m)$ in the atmosphere is given by the expression

$$\frac{\epsilon(m)}{\epsilon(m)_\odot} = (1-f) + \frac{f\epsilon(m)_{\rm ISM}}{\epsilon(m)_\odot} = (1-f) + fD_m\,, \qquad (1)$$

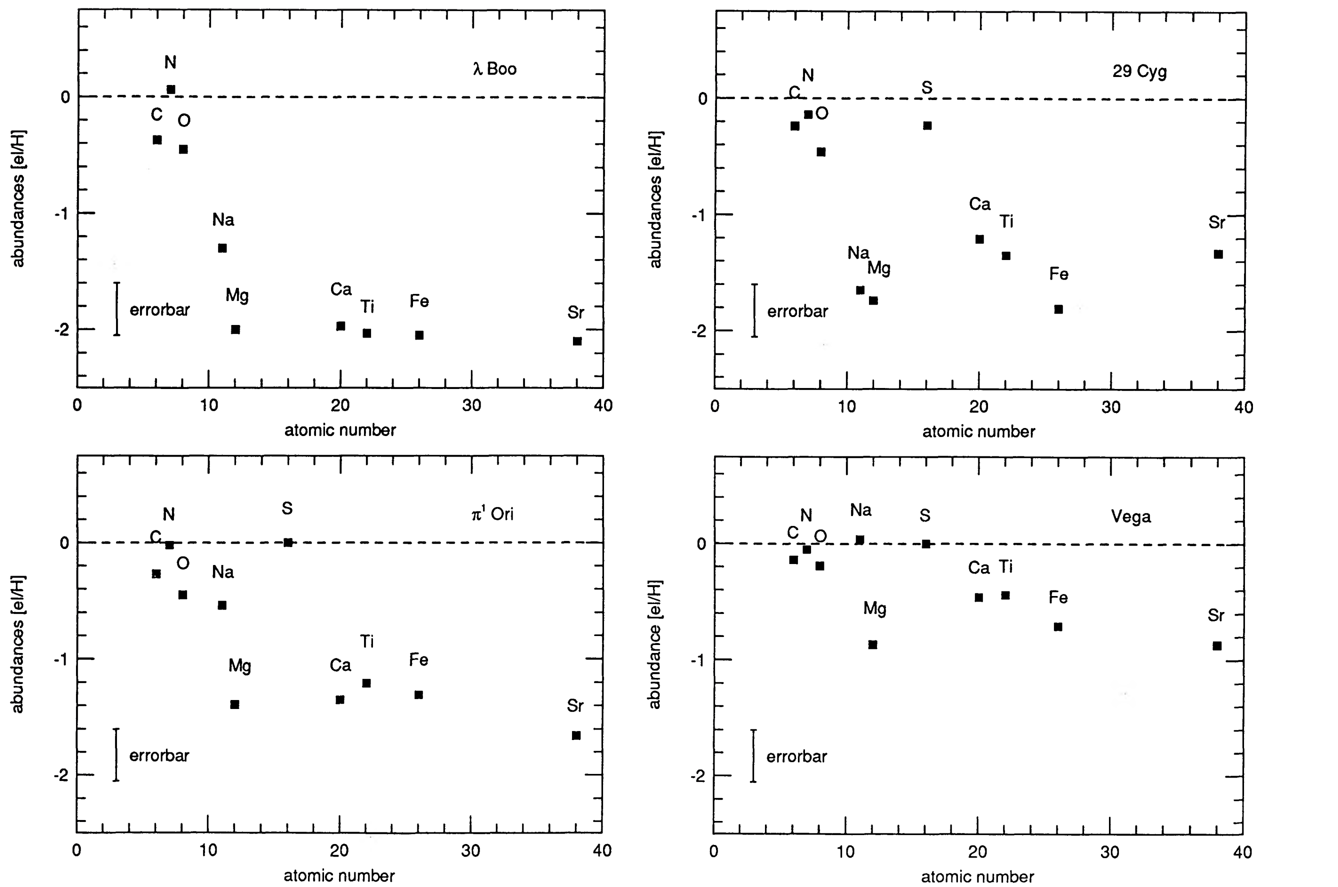

FIG. 4.—The chemical compositions of the λ Boo stars and Vega

where D_m is the depletion factor for element m in the interstellar gas.

Interstellar gas is highly depleted in many metals, but only slightly depleted in C, N, O, and S, and certain other elements. It was this signature of depletion that led us to the idea that the atmosphere of a λ Boo star consists of circumstellar gas mixed with a small amount of normal material. Equation (1) shows that the abundance of an element that is severely depleted in

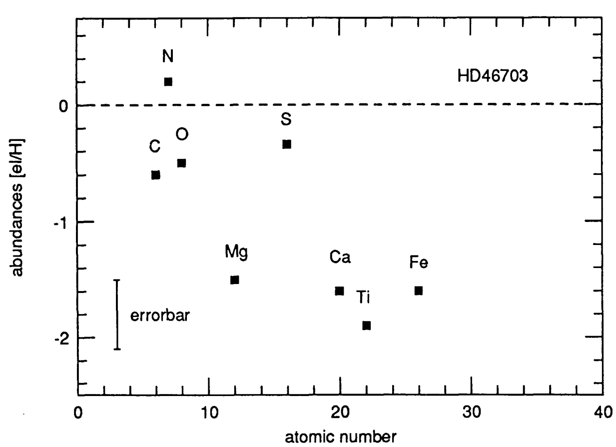

FIG. 5.—The chemical composition of HD 46703 (Luck and Bond 1984; Bond and Luck 1987).

the interstellar gas ($D_m \ll 1$) is

$$\frac{\epsilon(m)}{\epsilon(m_\odot)} \simeq (1-f)\,, \tag{2}$$

when $D_m < (1-f)$; i.e., elements satisfying this latter condition are deficient by a nearly constant factor in the stellar atmosphere. Moreover, the abundances $\epsilon(m)$ of elements for which $D_m \simeq 1$ are approximately solar independently of the mixing fraction f. Hence, the simple model explains why C, N, O, and S have similar and near-normal abundances, but the metals are underabundant by a nearly uniform factor that varies from star to star. Of course, when $D_m > (1-f)$, the relative abundances in the stellar atmosphere reflect the different depletion factors prevailing in the circumstellar gas; e.g., Ca is then underabundant relative to Fe.

Predicted compositions for mixtures of normal and circumstellar gas at $n(H_{tot}) = 3$ cm^{-3} are shown in Figure 6 for three values of the mixing fraction f (= 0.70, 0.95, and 0.99): $\bar{n}(H_{tot})$ is the average of the total H density [$= n(H) + 2n(H_2)$] along the line of sight. Depletion factors are taken from Jenkins (1987, 1988) or references cited therein. The depletions of the elements vary differently with the gas density, but the predicted relative abundances are a weak function of $n(H_{tot})$.

Inspection of Figures 4 and 6 show that the predicted compositions resemble the observed compositions of Vega and the three λ Boo stars. Sr cannot be considered because the interstellar abundance is unknown. For Vega, the observed com-

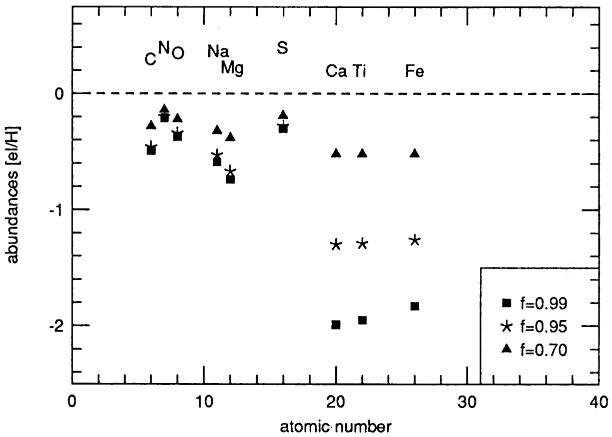

FIG. 6.—Predicted compositions for gaseous mixtures obtained by combining normal/solar material with interstellar gas. The mixing fraction f is defined in the text.

position is predicted to within about the expected uncertainties by the model corresponding to $f = 0.70$. Larger values of f are needed to match the compositions of the λ Boo stars: $f \simeq 0.95$ for π^1 Ori to $f = 0.99$ for λ Boo. For our sample of elements, the only obvious discrepancies between the observed and predicted abundances arise for Na and Mg; the predicted underabundances are much smaller than the observed values. We cannot propose a satisfactory resolution of these discrepancies. A non–LTE analysis of the Na I spectrum may show that the neutral atoms are overionized by the stellar radiation field and, hence, that our abundances are underestimates. The interstellar Na abundance is among the least well determined because the neutral atoms that contribute the Na D lines are a trace species and the Na^+ ion is not detectable. Initially, we adopted Morton's (1974) depletions for the ζ Oph cloud for which he gives $D_m \simeq 10^{-1.5}$ for Mg, a value that leads to predicted abundances in fair agreement with the observed compositions of the λ Boo stars. A revision of the gf-values for the weak Mg II interstellar lines (upon which the Mg abundance is based) has introduced a change in the D_m value [$= 10^{-0.7}$ for $\bar{n}(H_{tot}) = 3$] and, hence, a larger predicted abundance (as in Fig. 6).

Even if the comparisons of the observed and predicted abundances are taken as supporting our speculation that accreted circumstellar gas is the dominant constituent of the atmospheres of the λ Boo stars and Vega, several key questions remain to be addressed: Under what circumstances can circumstellar (or interstellar) gas be accreted with minimal accretion of the dust grains? What are the necessary conditions for the accreted gas to be the dominant component of the stellar atmosphere? Why is this form of accretion (apparently) restricted to a few peculiar stars?

Perhaps the association of circumstellar dust with the peculiar stars is a clue to answering such questions. The dust cloud or disks may have formed from interstellar grains in the star's parental cloud. Gravitational sedimentation of dust grains in interstellar clouds has been studied theoretically by Flannery and Krook (1978) and Harrison (1978)—see a brief review by Cassen and Boss (1988). It is evident that sedimentation is a slow process. It remains to be shown that dust-free gas was accreted efficiently by the star and remains largely unmixed with the stellar envelope that presumably has an essentially normal composition. Furthermore, the reasons why these particular stars develop a nebula and accrete gas but not grains remain unknown. It is, of course, possible that the process is quite common but rarely observable. The high rotational velocity of the λ Boo stars may be the result of the accretion process. As the rapidly rotating thin layer is mixed with the deeper layers, signs of rapid rotation and anomalous abundances disappear. One may offer a speculation on why A-type stars exhibit the effects of accretion. In lower mass stars, the deeper convective envelopes will quickly dilute the accreted material so that the abundance anomalies may never be remarkable. Higher mass stars may evolve too quickly for gas and dust to separate in the circumstellar nebula and grains driven off by radiation pressure then drag the gas out too.

There is a possibility that the similarity between the elemental depletions in the λ Boo stars and the interstellar gas arises from independent physical effects that are correlated with the same or related atomic properties. One such property is the first ionization potential: first (FIP) and second (SIP) ionization potentials are listed in Table 4. Snow (1975) noted that the interstellar depletion factors D_m are fairly well correlated with the FIP. Then, there is the observation that elemental abundances in the solar wind and solar photosphere differ by an amount dependent on the FIP. Although the association of the peculiar stars with circumstellar dust suggests a direct link between the circumstellar gas and the stellar photosphere, alternative interpretations should not be dismissed in haste.

The ratio of the elemental abundances in the extended solar atmosphere (solar corona, solar wind, and solar energetic particles) to the spectroscopically determined photospheric abundances is dependent on the ionization potential of the neutral atoms (the first ionization potential, FIP). For elements with FIP $\lesssim 10$ eV, the abundances relative to Si are identical in the extended atmosphere and photosphere, but elements with FIP > 11 eV are underabundant by a factor of 4 in the extended atmosphere (Cook, Stone, and Vogt 1984; Meyer 1985; Stone 1989). This dependence on the FIP is not understood, but it is supposed to indicate that fractionation takes place at the top of the photosphere where low FIP elements are largely singly ionized but high FIP elements are neutral. Acceleration and injection into the extended atmosphere are achieved more readily for ions than for neutrals. A similar process may be operating in the A stars. Of course, the fact that, in these stars, the neutral atoms are in a minority may modify the efficiency of the fractionation. In order to affect the photospheric composition, mixing between the photosphere and the deeper envelope must be suppressed. Figure 7 shows the abundances for λ Boo stars plotted against the FIP. Except for the Na abundance, the abundance-FIP correlation is similar in form to that shown for the Sun. A tighter correlation of a different shape is found using the second ionization potential (SIP): Na, an alkali atom, has the highest SIP of the sample, but the ordering of the other elements is essentially undisturbed. The correlations with the FIP or the SIP are as good or, perhaps, even tighter than that between the observed abundances and those predicted for mixtures of normal and circumstellar gas (Figs. 4 and 6). However, the latter is interpretable in terms of a very simple model and linked to the observed presence of dust around the stars.

V. CONCLUDING REMARKS

This study has extended the few previously published determinations of the chemical composition of λ Boo stars. In particular, the stars are metal-poor, but with approximately

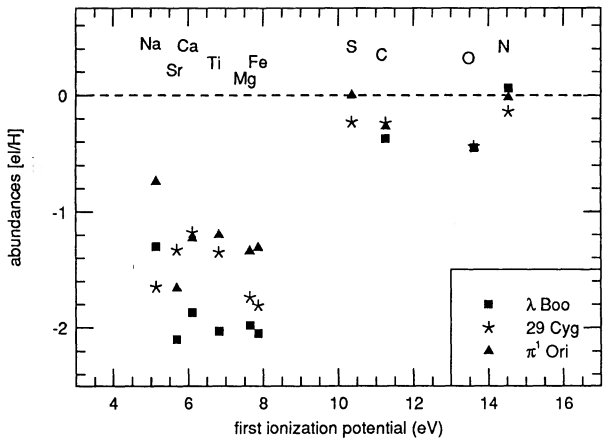

FIG. 7.—Abundances (relative to the Sun) for the three λ Boo stars vs. the first ionization potential (FIP) of the elements.

normal abundances of C, N, O, and S. The metal deficiencies span a range from [Fe/H] ~ -2 (λ Boo) through -1.3 (π^1 Ori) and to -0.7 if Vega is included as a mild slowly rotating λ Boo star.

The abundance anomalies are reminiscent of the depletions exhibited by the elements in the interstellar gas. This similarity and the presence of circumstellar material around at least some of the stars encouraged our speculation that the dominant component of the stellar atmospheres is accreted circumstellar (or interstellar) gas. Obvious tests of this speculation deserving attention include: the extension of the abundance analysis to other elements whose abundances may be predicted by our speculation (i.e., Zn should not share the underabundance of Fe); abundance analyses of the candidate λ Boo stars identified by Abt (1984) and Gray (1988) in order to determine the diversity of the abundance anomalies and, in particular, to test the prediction that, in the most extreme of the metal-poor stars, the relative abundances of the metals resulting from accretion of interstellar gas should resemble those in the interstellar gas; early-type stars (Aumann 1985; Sadakane and Nishida 1986) shown by the *IRAS* survey to possess circumstellar material (i.e., Vega-like stars) should be analyzed to check our prediction that their abundance anomalies may resemble those of the λ Boo stars and Vega.

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