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The stellar metallicity–giant planet connection

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ABSTRACT

The parent stars of the recently announced planetary system candidates are far from typical in terms of their chemical compositions. In this study we report on spectroscopic abundance analyses of υ And and τ Boo. Both stars are metal-rich relative to the Sun, with a mean [Fe/H] value near 0.25. These findings follow the trend set by two other planetary system candidates, ρ^1 55 Cnc and 51 Peg, which also display metallicities much higher than the average for nearby dwarfs. In addition, their companions share similar orbital characteristics. Given these observations, we propose that the current metallicities of these four stars are not representative of that of the original interstellar clouds from which they formed but, rather, are the result of self-pollution during the planet formation epoch early in their histories.

Key words: stars: abundances – stars: individual: HR 458 – stars: individual: HR 5185.

1 INTRODUCTION

Since the first announcement of a planet orbiting another solar-type star in 1995 October (Mayor & Queloz 1995), several additional candidates have been announced (Butler & Marcy 1996; Butler et al. 1996; Marcy & Butler 1996) at a rate averaging near one per month. Some effort has also gone into trying to determine the intrinsic properties of the parent stars (François et al. 1996; Henry et al. 1996; Perryman et al. 1996; Baliunas et al. 1997). Since we understand the physical properties of solar-type stars (F–G dwarfs) better than most other kinds of astronomical bodies, it should be possible to set some constraints on scenarios proposed to account for the companions by studying their parent stars. The metallicity of the interstellar cloud from which a star forms, for example, might have a significant effect on the planet formation process, since the initial build-up of planets is thought to occur through the accumulation of rocky planetesimals (Lissauer & Stewart 1993).

We have shown in a previous study (Gonzalez 1996) that the planetary system candidates ρ^1 55 Cnc and 51 Peg are probably super metal-rich (SMR) stars, which are defined as having [Fe/H] > 0.2 (Taylor 1996).[1] The companions of these two systems are similar in their orbital characteristics (periods near 4 d and masses near 0.6 M_J). These observations led us to propose a link between the existence of a giant planet with a small orbit and the high metallicity of its parent star, which makes use of the hypothesis of Lin, Bodenheimer & Richardson (1996). This model involves the inward migration of a gas giant during the planet formation epoch while the accretion disc is still present. As the planet spirals inward in the disc, the material between it and the star presumably falls on to the latter. If the star accretes material deficient in H and He, then the metallicity of the parent star might be enhanced significantly.

In order to gather additional data with the goal of testing this hypothesis further, we have performed abundance analyses on the parent stars of two additional planetary system candidates, υ And (=HR 458) and τ Boo (=HR 5185). Like the 51 Peg and ρ^1 55 Cnc systems, the planetary companions of these two stars also have very short orbital periods (Butler et al. 1996).

2 OBSERVATIONS

Spectra of υ And and τ Boo were obtained on 1996 July 18/19 with the echelle spectrograph at the coudé focus of the 2.7-m telescope at the McDonald Observatory. The spectra were imaged on to a 2048 × 2048 Tektronix CCD giving spectral coverage from 3900 to 10 200 Å, with gaps in the spectra beyond 5600 Å, and a resolving power ($\lambda/\Delta\lambda$) of ~60 000 (instrument described by Tull et al. 1995). Spectra of the Sun, reflected off the asteroid Vesta, and of a hot star with a high rotation velocity were also obtained during the same run.

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[1]We use the standard spectroscopic notation, $[X/Y] \equiv \log_{10}(N_X/N_Y)_{star} - \log_{10}(N_X/N_Y)_{Sun}$ and $\log \varepsilon(X) \equiv \log_{10}(N_X/N_H) + 12$, for number density abundances of elements X and Y.

The spectra were reduced using the standard data analysis programs available in the ECHELLE package of NOAO IRAF. The wavelength scale was calibrated with a Th-Ar hollow-cathode spectrum obtained on the same night; the standard deviation of the calibration is less than 2 mÅ. The spectrum of the hot star was used to divide out telluric lines in spectral regions containing photospheric lines used in the abundance analysis. The S/N ratios in the 5800–7000 Å spectral region, after division by the hot star spectrum, are typically 600–800 for the spectra of υ And and τ Boo, and 400–500 for the spectrum of Vesta. Most of the equivalent widths (EWs) have been determined by integrating the area under a line, but a few have been estimated by fitting Gaussian functions; the typical uncertainty in an EW measurement is less than 1 mÅ. We present sample spectra of υ And and τ Boo in Fig. 1.

3 ANALYSIS

3.1 Method

In order to achieve the highest accuracy possible, we employ a differential abundance analysis relative to the Sun; such an analysis minimizes systematic errors, especially for solar-type stars. It is similar to the studies of Gonzalez (1996) and Gonzalez & Lambert (1996), who scaled laboratory based *gf*-values using EWs measured on the solar spectrum. We do not use their *gf*-values, since our data were obtained with a different instrument. Instead, we first estimate the solar iron abundance and depth-independent microturbulence parameter, ξ_t, using the laboratory Fe I *gf*-values given in Lambert et al. (1996) and the EW measurements on our spectrum of Vesta. Next, using this solution, the *gf*-values of

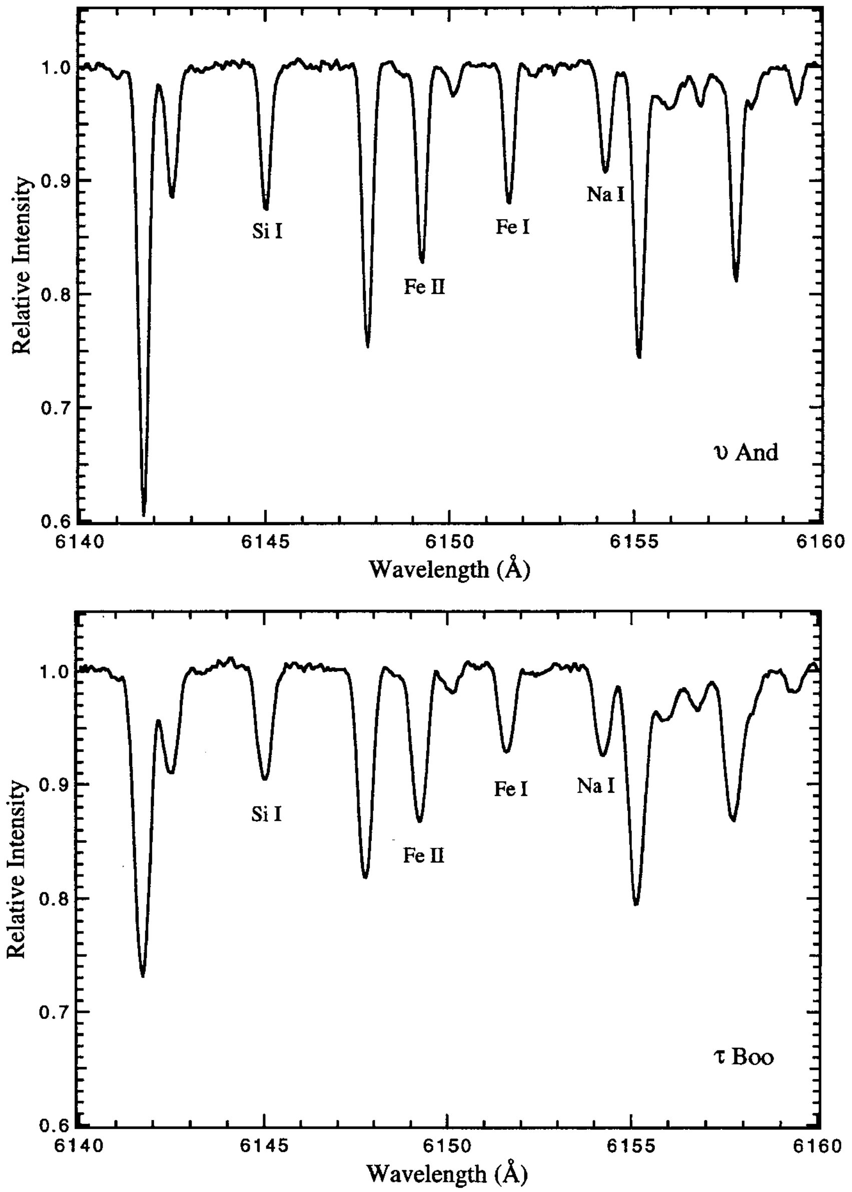

Figure 1. Sample spectra of υ And and τ Boo. Only those lines used in the abundance analyses are labelled. The lines in the τ Boo spectrum appear weaker because they are broader. The lines are actually quite similar in strength in the two spectra.

other spectral lines are estimated from an inverted solar analysis. This involves adjusting the *gf*-value of a given transition so that it gives the desired abundance (in the present work we adopt the photospheric solar abundances of Anders & Grevesse 1989 for all elements except Li and Fe).

The abundance analyses make use of the LTE plane-parallel model atmospheres of Kurucz (1993) with an updated version of the line abundance analysis code MOOG (Sneden 1973). Although the spectral coverage is extensive and potentially thousands of absorption lines are available in each spectrum, we have been very restrictive in the selection of lines for analysis. The same set of lines are measured in all three spectra, and only Fe I lines unblended in all three spectra are used in the analysis. Given that the linewidths in the spectra of υ And and τ Boo are about 1.9 and 3.0 times those in the Vesta spectrum, respectively, lines that are unblended in the Vesta spectrum might be blended in the other two spectra due to their much greater line broadening. Hence the number of lines available for analysis is set primarily by the τ Boo spectrum. These selection criteria have been relaxed slightly for the other atomic species.

3.2 Results

Based on the Fe I line analysis of the Vesta spectrum, the value of $\log \varepsilon$(Fe I) is 7.47, and ξ_t is 1.0 km s^{-1}; these estimates are consistent with the results of other similar studies (Gonzalez & Lambert 1996). The physical parameters of υ And and τ Boo have been estimated using 20 Fe I and three Fe II lines. The solutions ($T_{\rm eff}$, $\log g$, ξ_t) are 6250 ± 100 K, 4.3 ± 0.1 (cgs), 1.4 ± 0.1 km s^{-1} and

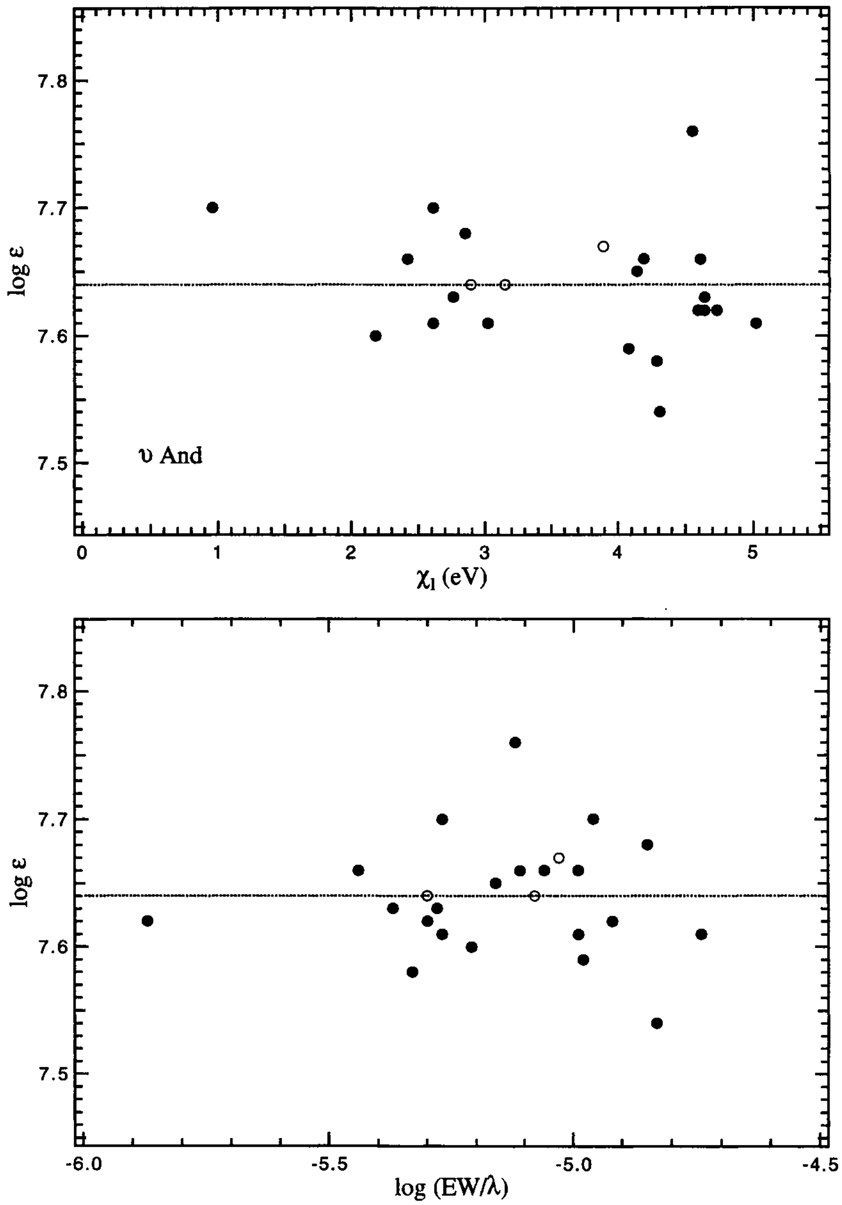

Figure 2. The Fe I (filled circles) and Fe II (open circles) individual line abundances of υ And, calculated using the parameters $T_{\rm eff} = 6250$ K, $\log g = 4.3$ (cgs), $\xi_t = 1.4$ km s^{-1} and [Me/H] = 0.20.

6600 ± 100 K, 4.5 ± 0.15 (cgs), 1.6 ± 0.1 km s^{-1} for υ And and τ Boo, respectively; the corresponding values of $\log \varepsilon$(Fe) are 7.64 ± 0.08 and 7.81 ± 0.09 (or equivalently, [Fe/H]$=0.17 \pm 0.08$ and 0.34 ± 0.09). The individual line abundance estimates plotted against χ_1 and $\log(\mathrm{EW}/\lambda)$ are presented in Figs 2 and 3 for υ And and τ Boo, respectively. The Fe I line at 6498 Å is blended with a Ca I line in the spectra of υ And and τ Boo (but not in the Sun), so the EW estimates for this line are less accurate than for the other Fe I lines; we retained it in the analysis since it improves the leverage in determining T_{eff}. The uncertainties in the abundance estimates have been calculated in the same way as Gonzalez (1996) and Gonzalez & Lambert (1996). The lithium abundance has been estimated by synthesizing the 6700 to 6710 Å spectral region, which includes the Li I doublet at 6707.8 Å. The individual line abundance estimates are listed in Table 1, and the mean abundances are listed in Table 2. Several elements are represented by only one line in the analysis; however, given the high quality of the selected lines, the estimated abundances of these elements, especially Na, S, Sc, Cr and Ni, should be considered quite reliable. The lines of N I, Mg I and Zn I are of slightly lower quality due to the presence of telluric lines or crowding. For species represented by more than one line, C I, Al I, Si I, Ca I, Ti I and Fe I, II, the abundances from individual lines of the same element show very small scatter.

We have also obtained spectra of τ Boo on 1996 June 25–27 with the 2.1-m telescope at McDonald Observatory and analysed them using the same procedure and linelist as Gonzalez (1996). The estimated values of T_{eff}, $\log g$, ξ_t and [Fe/H] are 6750 ± 150, $\log g = 4.6 \pm 0.15$ (cgs), 1.5 ± 0.2 km s^{-1} and 0.53 ± 0.14, respectively. The abundance estimates of the other elements are also consistent with the present results.

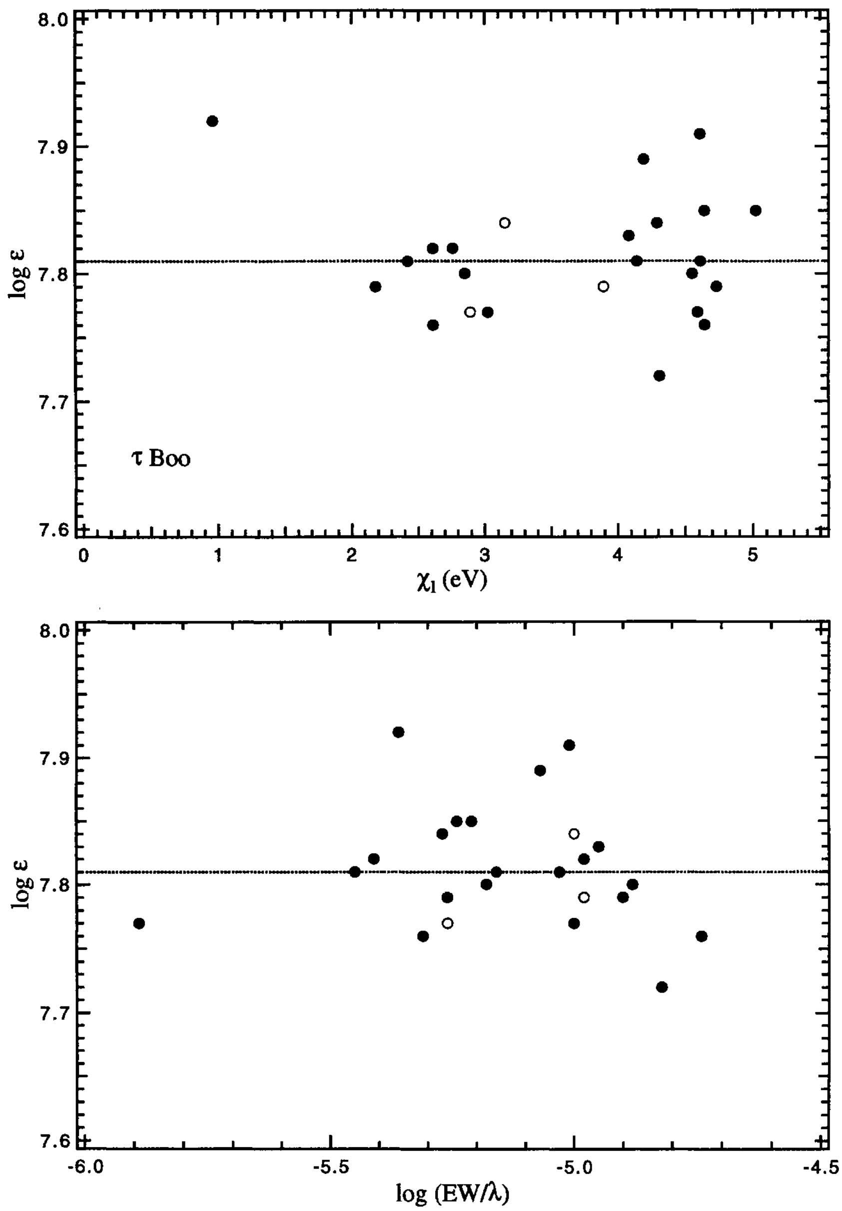

Figure 3. Same as Fig. 2, but for τ Boo with the following parameters: $T_{\mathrm{eff}} = 6600$ K, $\log g = 4.5$ (cgs), $\xi_t = 1.6$ km s^{-1} and [Me/H]$=0.30$.

Table 1. Measured equivalent widths and derived abundances for the programme stars.†

Species λ (Å)	EW (Vesta) mÅ	log *gf*	EW, log ε υ And	EW, log ε τ Boo
Li I				
6707.80	SS	SS	—, 2.42	—, 1.60
C I				
7108.92	11.3	-1.18	13.7, 8.28	22.8, 8.47
7115.62	28.1	-0.65	40.7, 8.41	53.0, 8.49
7116.96	21.9	-0.81	—, —	52.6, 8.65
N I				
7468.27	5.2	-0.02	9.2, 7.90	16.7, 8.06
Na I				
6154.23	39.8	-1.58	32.8, 6.46	34.6, 6.63
Mg I				
5711.10	112.0	-1.71	112.0, 7.65	98.8, 7.82
Al I				
7835.29	44.4	-0.71	44.5, 6.65	42.6, 6.74
7836.11	55.6	-0.57	61.3, 6.74	65.6, 6.89
Si I				
5793.06	45.1	-1.91	49.2, 7.75	56.8, 8.00
6125.41	32.6	-1.54	36.0, 7.73	43.5, 7.96
6145.41	39.7	-1.42	45.9, 7.77	49.1, 7.92
S I				
6052.65	12.3	-0.44	25.0, 7.34	38.2, 7.54
Ca I				
5867.55	24.1	-1.62	23.5, 6.60	21.6, 6.72
6166.44	71.7	-1.13	64.2, 6.49	64.6, 6.66
Sc II				
6604.58	36.2	-1.17	46.5, 3.27	44.8, 3.37
Ti I				
5965.84	32.1	-0.38	22.1, 5.13	19.8, 5.32
6126.21	22.7	-1.41	14.3, 5.18	13.4, 5.44
V I				
5727.06	40.2	0.00	23.4, 4.06	16.3, 4.16
Cr I				
5787.93	47.6	-0.11	41.3, 5.82	48.8, 6.13
Fe I				
5044.21*	74.9	-2.04	70.8, 7.68	67.1, 7.80
5806.73*	54.4	-0.90	51.0, 7.66	56.4, 7.91
5852.22*	41.3	-1.18	44.7, 7.76	38.3, 7.80
5855.09	23.0	-1.52	21.1, 7.66	20.7, 7.81
5856.10	34.4	-1.56	27.6, 7.58	31.4, 7.84
6027.05*	64.9	-1.09	63.2, 7.59	61.8, 7.83
6056.01*	75.7	-0.40	73.5, 7.62	76.0, 7.79
6065.48*	119.5	-2.04	111.0, 7.61	110.1, 7.76
6089.57	36.6	-0.86	32.5, 7.61	37.9, 7.85
6151.62*	51.1	-3.29	37.6, 7.60	33.6, 7.79
6165.36*	45.9	-1.47	43.0, 7.65	42.2, 7.81
6200.31*	74.6	-2.44	68.5, 7.70	64.9, 7.82
6380.74*	54.1	-1.32	50.0, 7.66	53.7, 7.89
6498.95	47.1	-4.62	34.8, 7.70	28.5, 7.92
6591.33	10.6	-1.98	8.9, 7.62	8.5, 7.77
6703.58	38.0	-3.01	28.7, 7.63	26.0, 7.82
6750.15*	76.0	-2.62	68.5, 7.66	63.6, 7.81
6752.71*	37.1	-1.20	33.8, 7.62	33.0, 7.76
6820.37*	41.7	-1.17	36.2, 7.63	39.6, 7.85
7583.80	85.2	-1.90	78.2, 7.61	76.4, 7.77
7586.03	130.8	-0.18	111.3, 7.54	114.8, 7.72
Fe II				
5991.38	33.4	-3.48	50.3, 7.64	59.3, 7.84
6149.25	36.8	-2.70	57.3, 7.67	64.9, 7.79
6369.45	19.6	-4.11	31.8, 7.64	34.8, 7.77
Ni I				
6767.76	80.9	-2.09	72.8, 6.34	69.9, 6.52
Zn I				
4722.15	74.2	-0.26	80.4, 4.75	76.6, 4.77

†The *gf*-values of the Fe I lines marked with an asterisk are from Lambert et al. (1996). The *gf*-values of the other lines have been determined from an inverted solar analysis using the EW values measured on the spectrum of Vesta. SS is an abbreviation for spectrum synthesis.

3.3 Photometric estimates

We can also derive estimates of the physical parameters of the programme stars from Strömgren *uvbyβ* narrow-band photometry. Hauck & Mermilliod (1990) give mean values of the Strömgren indices for these stars; Olsen (1994) gives more recent values, which are within a few mmag of the mean values. Using the $T_{\rm eff}$ calibrations of Saxner & Hammarbäck (1985), we estimate $T_{\rm eff}$ to be 6168 and 6405 K for υ And and τ Boo, respectively; the corresponding values of [Fe/H] using equation (8) of Nissen (1988) are 0.00 and 0.18 dex. Reddening corrections have not been applied, since these stars are nearby. The Bright Star Catalog (Hoffleit & Jaschek 1982) notes the presence of an M2 star about 5 arcsec from τ Boo. However, given that it is about 7.5 mag fainter than the primary, it should not influence the photometric colours significantly.

Table 2. Final adopted abundances relative to the Sun for υ And and τ Boo.

Atomic Species	log ε⊙	N	υ And [X/H]	τ Boo [X/H]
Li I	1.06	1	1.36 ± 0.06	0.54 ± 0.30
C I	8.56	5	-0.22 ± 0.09	-0.03 ± 0.08
N I	8.05	1	-0.15 ± 0.05	0.01 ± 0.07
Na I	6.33	1	0.13 ± 0.06	0.30 ± 0.07
Mg I	7.58	1	0.07 ± 0.08	0.24 ± 0.09
Al I	6.47	2	0.23 ± 0.05	0.35 ± 0.08
Si I	7.55	3	0.20 ± 0.05	0.41 ± 0.07
S I	7.21	1	0.13 ± 0.05	0.33 ± 0.07
Ca I	6.36	2	0.19 ± 0.08	0.33 ± 0.06
Sc II	3.10	1	0.17 ± 0.06	0.27 ± 0.06
Ti I	4.99	2	0.17 ± 0.05	0.39 ± 0.07
Cr I	5.67	1	0.15 ± 0.06	0.46 ± 0.07
Fe I, II	7.47	21, 3	0.17 ± 0.08	0.34 ± 0.09
Ni I	6.25	1	0.09 ± 0.08	0.27 ± 0.10
Zn I	4.60	1	0.15 ± 0.08	0.17 ± 0.09

3.4 Other physical characteristics

The absolute magnitudes, M_v, of υ And and τ Boo are easily derived from their parallaxes (Lutz & Kelker 1973; van Altena, Lee & Hoffleit 1995) and visual magnitudes; we calculate $M_v = 2.80 \pm 0.15$ and 3.11 ± 0.13, respectively. Combining these estimates of M_v and the other derived physical parameters (from Section 3.2) with the VandenBerg (1985) theoretical stellar evolutionary tracks, it is possible to estimate their masses and ages. Doing so and applying corrections to the model tracks to give the correct age of the Sun (see Gonzalez 1996), and extrapolating them to the higher metallicities of υ And and τ Boo, we estimate the masses to be 1.34 ± 0.08 and $1.55 \pm 0.06\,{\rm M}_\odot$ and the ages to be 4 ± 1 and 1.25 ± 0.25 Gyr, respectively.

We have measured the projected rotational velocity, $v \sin i$, and the macroturbulent velocity, $\zeta_{\rm RT}$, for each star using the same procedure as Gonzalez; the results for υ And and τ Boo are $v \sin i = 9.0 \pm 0.5$ and 14.5 ± 0.5 km s^{-1}, and $\zeta_{\rm RT} = 6.0 \pm 0.5$ and 8.5 ± 0.5 km s^{-1}, respectively. These values of $v \sin i$ are consistent with the young ages derived above, and the estimates of $\zeta_{\rm RT}$ are typical of F dwarfs with temperatures near those estimated in Section 3.2 (Gray 1992, fig. 18.9). Combining our $v \sin i$ estimates with the mass functions determined by Butler et al. (1996) and the rotation periods quoted by them, we calculate that the masses of υ And B and τ Boo B are $0.77^{+0.17}_{-0.04}$ and $6.7^{+6.2}_{-1.9}\,{\rm M_J}$, respectively.

4 DISCUSSION

4.1 Comparison with previous work

Previous spectroscopic abundance estimates for υ And and τ Boo can be found in the literature; Cayrel de Strobel et al. (1992) list three estimates of [Fe/H] for υ And ranging from -0.11 to -0.23 dex, and four estimates for τ Boo ranging from 0.00 to 0.30 dex. In addition, Edvardsson et al. (1993) derived [Fe/H] = 0.09 for υ And from a spectroscopic analysis supplemented with photometric data. The study of Boesgaard & Lavery (1986) included both stars, obtaining [Fe/H] = 0.06 for υ And and [Fe/H] = 0.30 for τ Boo; their results are based on high-resolution, high S/N ratio spectra, but the temperatures are derived from photometry, and only nine Fe I lines are used in their analysis. Also, their use of a mean ξ_t value, 1.8 km s^{-1}, for all their programme stars introduces systematic errors into their abundance estimates; use of a smaller value brings their metallicity estimates closer to our results. We consider our estimates to be more reliable than those determined by previous studies due to our use of very high-quality spectra and our complete reliance on spectroscopy to determine the fundamental stellar parameters. However, taking into account the uncertainties of all estimates, our results are generally consistent with previous studies.

4.2 Words of caution

As we showed in Section 3.3, the temperatures estimated from photometry are lower than those derived from spectroscopy; the difference is minor for u And, but it is about 200 K for τ Boo. This might be due to the scarcity of SMR standard calibrators and the consequent relatively poor calibration of photometric indices for metal-rich stars. Our estimates of $T_{\rm eff}$ are unlikely to be in error by more than ± 150 K and, as shown in Section 3.4, they are consistent with our estimates of $\zeta_{\rm RT}$, which is a simple function of spectral type and luminosity class.

Petit (1990) lists τ Boo as a possible δ Scuti variable and quotes a visual apparent magnitude ranging from 4.46 to 4.52. Further evidence of its variability is given by Butler et al. (1996), who report that τ Boo exhibited erratic velocity variations during portions of their observing runs; the velocities deviate from a Keplerian fit by nearly 100 m s^{-1}. Such variations, if due to some kind of pulsation, are not likely to account for temperature variations near 200 K, but this possibility cannot be dismissed until the nature of the variations is better understood. A careful study of variations in the line-profile shapes similar to that of Hatzes, Cochran & Johns-Krull (1996) would be very helpful in interpreting the velocity variations of the photosphere of τ Boo.

4.3 Origin of high metallicities

The high estimates of [Fe/H] for υ And, τ Boo, ρ^1 55 Cnc and 51 Peg are exceptional when compared to the metallicity distribution of nearby stars; two of them definitely belong to the rare class of SMR stars, while υ And is slightly below the SMR limit and 51 Peg is near the limit. About 2 per cent of the nearby G dwarfs have [Fe/H] > 0.2, as determined from photometry (Rocha-Pinto & Maciel 1996); interestingly, this is very close to the frequency of occurrence of 51-Peg-like systems among the stars searched for radial velocity variations (Butler et al. 1996). The high incidence of metal-rich stars in our sample, which have been selected for study on the basis of radial velocity variations due to the presence of unseen companions, suggests that the metal-richness is somehow linked to the presence of planets – assuming, of course, that the companions are planets. What is the nature of this link? Is the high metallicity a prerequisite for the formation of giant planets, or does the planet formation process alter the stellar surface abundances, or is it a combination?

Given the currently popular mechanism of giant planet formation involving first the accretion of planetesimals until a mass of 10–15 $M_\oplus$ is reached and then the capture of H and He gas (Podolak, Hubbard & Pollack 1993), one might expect that a certain minimum metallicity is required for this process. The existence of Jupiter in the Solar system implies that this limit in [Fe/H], if it exists, is less than or equal to 0.0. By itself, a high metallicity does not explain the very small orbital radii of the companions of these four stars. However, if we assume that these companions are gas giants that were formed outside their current orbits and then migrated inward (Lin et al. 1996), then most of the disc material between the planet and the star was probably accreted on to the latter. This might have increased the envelope metallicity of the star significantly; the addition of 20 $M_\oplus$ of chondritic material to the Sun in its early history would have increased [Fe/H] in the convective envelope by about 0.11 dex (Gonzalez 1996). The accreted material is mixed throughout the convective region of the envelope, which comprises only about 3 per cent of the mass of the star near the main sequence. The envelope convective region is even smaller for F dwarfs. Hence, all else being equal, the accretion of a given quantity of gas-poor material on to the surface of τ Boo should have a more pronounced effect on its surface abundances than it would on the other three systems. The higher metallicity of τ Boo is consistent with this scenario. However, ρ^1 55 Cnc, which is cooler than 51 Peg, has a similar metallicity. In the light of the self-enrichment scenario, this is surprising considering that the convective region of ρ^1 55 Cnc is considerably larger than that of 51 Peg. If ρ^1 55 Cnc was originally enriched as much as τ Boo, then this might explain its currently high metallicity.

Another, less direct, piece of evidence in favour of the self-enrichment scenario is the great evolutionary age of ρ^1 55 Cnc, about 14 Gyr (Gonzalez 1996). There is a general age–metallicity relation among disc stars in the Milky Way, where the older stars are more metal-poor. The relation has considerable scatter; the observed range of [Fe/H] for 14-Gyr-old stars is roughly -0.4 to -0.9 (Edvardsson et al. 1993). Our estimate of [Fe/H] for ρ^1 55 Cnc (0.24 ± 0.09) is outside this range, implying that it was originally more metal-poor. However, a possible problem with the evolutionary age estimates for these SMR stars is the assumption of homogeneous composition in the stellar models, which is invalidated by the self-enrichment scenario. New heterogeneous models will need to be developed, where the envelope metallicity is greater than that of the interior.

Finally, we note that there are no definite SMR K giant stars, while there are eight (counting τ Boo and ρ^1 55 Cnc) SMR dwarfs (Taylor 1996). This fits the self-enrichment scenario for the following reason: as a self-enriched dwarf leaves the main sequence and ascends first the subgiant and then the giant branch, the outer convection zone deepens, diluting the metal-enriched outer layers with the interior. There is a potentially complicating factor, though; before the red giant phase is reached, the Jupiter-mass companions of the 51-Peg-like systems will be engulfed by their parent stars. If the companions are gas giants, then they will not alter the surface metallicity significantly (except for lithium); if they are rocky planets, then the surface metallicity will increase.

The fact that 47 UMa B has not experienced a significant planetary migration process and that 47 UMa is less metal-rich than the 51 Peg-like systems implies that a minimum metallicity is required to trigger planet migration.

4.4 Is the Sun a metal-enriched star?

Gies & Lambert (1992) and Snow & Witt (1996) have discussed evidence indicating that the Sun's photospheric composition is not representative of the nearby Galactic neighbourhood. Specifically, the solar abundances appear to be enhanced relative to B stars, the Orion nebula, and young field F and G stars. Even F–G dwarfs of similar age and Galactic orbits to the Sun have a lower mean metallicity: $\langle[\mathrm{Fe/H}]\rangle = -0.14 \pm 0.27$ (Edvardsson et al. 1993); even though the Sun is not an extreme point in this sample, it is near the upper envelope of the distribution. To illustrate this point, we show in Fig. 4 the mean abundances of B and nearby F–G dwarf stars relative to the solar photospheric abundances plotted against the condensation temperatures.[2] Given that the mean metallicity in the Galactic disc increases with time, it is surprising that the Sun is more metal-rich than its younger neighbours. The explanation usually offered is that the Galactic gas is poorly mixed, and that the Solar system formed in a more metal-rich clump.

We offer an alternative explanation: the Sun accreted H- and He-poor material during the planet formation epoch, which resulted in the increase of the abundances of the metals in its convective envelope. Like 47 UMa, the Sun might have become self-enriched via the Poynting–Robertson effect or comets. The magnitude of the effect would have to be such that the surface abundances are increased by at least 0.1–0.2 dex. Also, if some fractionation occurred during the accretion process, or if previously fractionated material was accreted, then one would expect a correlation of the solar photospheric abundances of the elements with their condensation temperatures. To test this hypothesis, we have plotted the differences between the solar photospheric and meteoritic abundances versus the condensation temperatures (Fig. 5) using the best available data (Grevesse, Noels & Sauval 1996). Although there is a hint of a correlation, the data are not yet of sufficient quality to say one way or the other. It would be especially helpful to reduce the uncertainties in the solar photospheric abundance estimates of sulphur and scandium, which have low and very high condensation temperatures, respectively. Also, the abundance difference for manganese is several times the quoted uncertainties; the reason for this is not clear. It should also be noted that the condensation temperature depends on the gas pressure; if the temperatures are calculated with the wrong pressure, then this will result in greater scatter in the relation. If the fractionation occurred in the Sun but not in other stars, then there should be correlations in the two diagrams presented in Fig. 4. Again the data do not allow us to reach a firm conclusion.

The Sun might not be unique among the sample of F–G dwarfs quoted above in experiencing a self-enrichment process. If it is not a rare occurrence among solar metallicity F–G dwarfs, then it is not possible to determine the magnitude of enrichment in the Sun compared to this sample of stars. If it is a stochastic process occurring in a fraction of F–G dwarfs, then this might account for the spread in metallicity among stars of the same age and location in the Galaxy. There is another anomaly: as is evident in Fig. 4, the mean metallicity among B stars is 0.1–0.2 dex less than it is among solar-age F–G dwarfs. Whether or not this difference is observationally significant is not clear at this time.

4.5 Implications of self-enrichment

There are a number of important implications resulting from the self-enrichment scenarios, particularly in the area of stellar structure and evolution, and possibly also Galactic chemical evolution; the basic assumptions that the metallicity of the envelope of a star is the same as that of its interior and also the same as the interstellar cloud from which it formed might need to be reconsidered. An error in the interior metallicity of a star will lead to an error in the predicted central temperature and, hence, its luminosity. This might have a significant bearing on the solar neutrino problem. Bahcall & Ulrich (1988) have calculated neutrino fluxes for solar models with metal-poor interiors; they estimate that a reduction of the metallicity in the layers below the convection zone by 1 dex can solve the solar neutrino problem. While it is unlikely that the amount of self-enrichment in the Sun is more than a few tenths of a dex, at least the effect is in the right direction.

The metallicity distribution of the nearby dwarfs, which has been very important in chemical evolution studies (Rocha-Pinto & Maciel 1996), might require a correction so that it accurately reflects the composition of the ISM at the time a star formed. If only the Lin et al. (1996) mechanism is responsible for self-enrichment, then only a small percentage of stars will require a correction, and the effect on the metallicity distribution will probably be minor. However, if an additional mechanism is in operation, such as Poynting–Robertson and comet accretion, then a greater number of stars would be affected, and the metallicity distribution would require a more significant correction. This may even go part way in solving the so-called G dwarf problem (Francois, Vangioni-Flam & Audouze 1990), where there is a deficit of metal-poor nearby G dwarfs relative to simple Galactic chemical evolution models.

[2]The condensation temperature of a given element is that temperature in a gaseous environment whereby half the atoms of that element condense out of the gas phase, usually on to grains. The condensation temperatures used in Figs 4 and 5 are from the first column of appendix G of Wasson (1985), corresponding to a gas pressure of 10^{-4} atm. The gas-phase abundances in the ISM correlate strongly with condensation temperature, such that those elements with the highest condensation temperatures have the lowest gas-phase abundances.

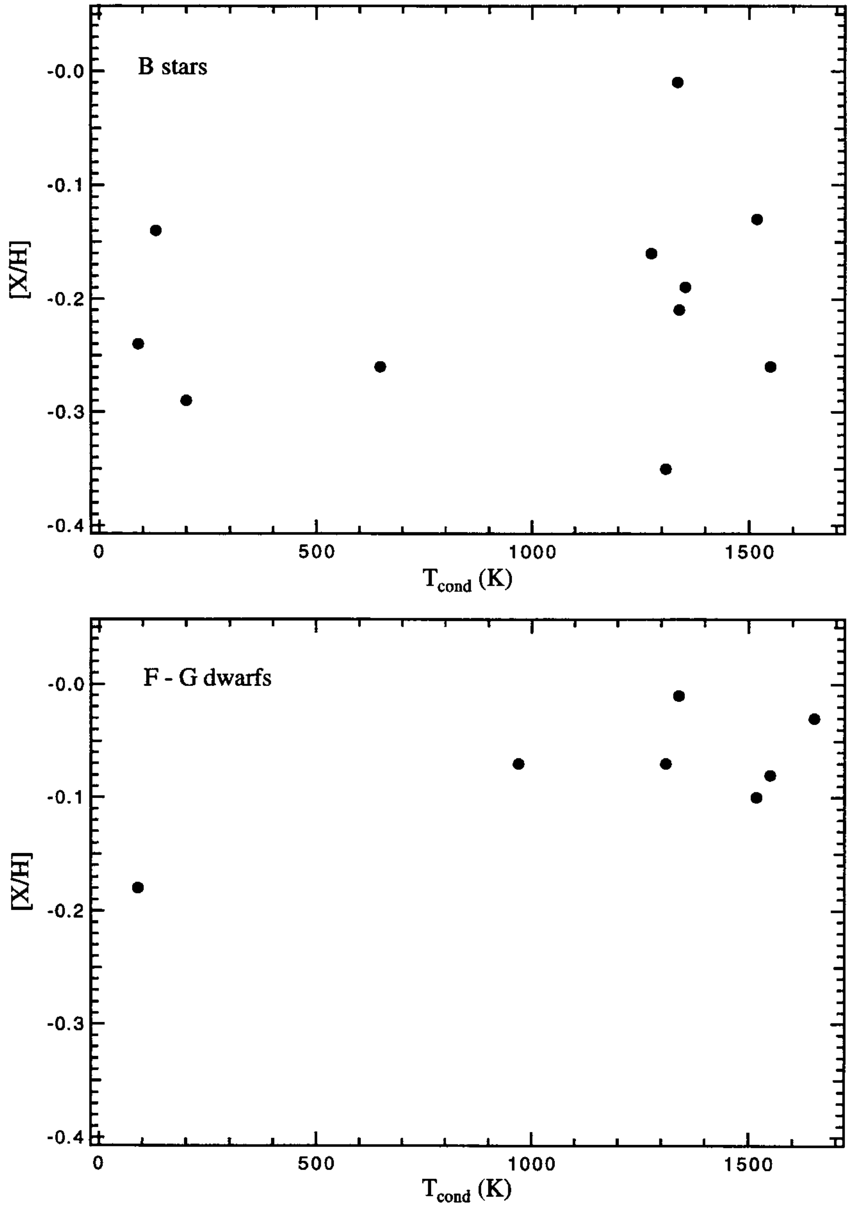

Figure 4. Mean abundances of B stars relative to the Sun (using the data from Grevesse, Noels & Sauval 1996 and Snow & Witt 1996) are plotted against the condensation temperature of each element (from Wasson 1985) in the first panel. The mean relative abundances of the nearby F–G dwarfs sharing the Sun's Galactic orbit (using the data of Edvardsson et al. 1993 and Tomkin et al. 1995) are shown in the second panel.

4.6 Selection effects

We should not dismiss the possibility that the apparent correlation of orbital parameters with the metallicity of the parent star among the 51-Peg-like systems is due to a selection effect. For example, a higher metallicity might lead to the formation of a more massive disc, which might be more likely to cause planet migration. Systems with massive planets in tight orbits are easier to detect. Hence, based on these assumptions, a more metal-rich system is more likely to exhibit radial velocity variations. However, this argument is weakened by the fact that the present radial velocity searches are able to detect other types of systems, such as 47 UMa. Probably the only definitive way to choose between the self-enrichment scenario and a selection effect is to examine planetary system candidates in a wide-separation binary or preferably in an open cluster, where the metallicities of stars with radial velocity variations can be compared to the cluster mean metallicity.

5 CONCLUSIONS

We have performed precise differential abundance analyses on υ And and τ Boo using high-resolution very high S/N spectra. The results indicate that τ Boo is a SMR star, and that υ And is slightly below the SMR limit. Combining these results with those of our previous study demonstrates that υ And, τ Boo, ρ^1 55 Cnc and 51 Peg are all members of a new

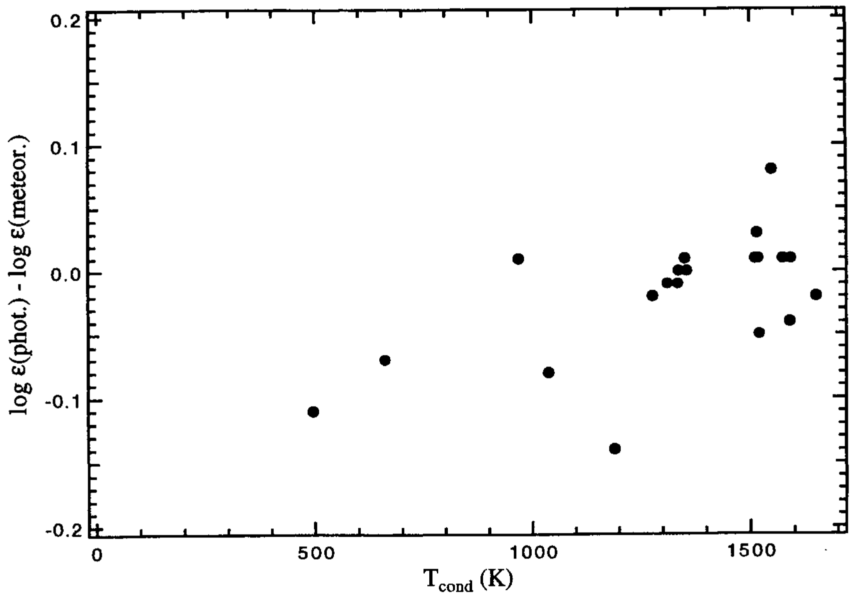

Figure 5. The differences between the solar photospheric and the meteoritic abundances (from Grevesse, Noels & Sauval 1996) are plotted against the condensation temperature (from Wasson 1985). Only those elements with standard deviations less than 0.1 dex in both the photospheric and meteoritic estimates, and with known condensation temperatures, have been included in the plot.

astronomical category. These systems share the following characteristics: metal-rich parent star and an approximately Jupiter-mass companion in a very small nearly circular orbit. The scenario of Lin et al. (1996), whereby a gas giant spirals inward and causes disc material to be accreted by the parent star during the early history of a stellar system, is the best explanation at this time for the existence of these systems. Less efficient self-enrichment mechanisms might have also operated in the 47 UMa and Solar systems.

The following suggestions for observers are the next logical steps to pursue in testing the self-enrichment hypothesis.

(1) An effort should be made to search for radial velocity variations among the known SMR stars. Such an effort is not difficult, given their scarcity.

(2) Spectroscopic abundance anlayses of the stars in a wide-separation binary system where one member is a planetary system candidate would tell us if self-enrichment has occurred. Optimally, both stars would be of spectral type F or G, which are easiest to analyse; we note that ρ^1 55 Cnc (Duquennoy & Mayor 1991) and τ Boo have M-type companions, but such stars have very complex spectra.

(3) Even more desirable would be the detection of planetary-mass companions in nearby Galactic star clusters, where small variations in metallicity among the cluster members can be studied with a statistically significant sample.

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Recently, a planetary-mass companion has been discovered to orbit 16 Cyg B, which is a member of a wide-separation binary system. Preliminary spectroscopic abundance analyses of these two stars indicates that they have the same metallicities to within 0.1 dex.

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