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The solar abundance of iron and the photospheric model[*]

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Abstract. Numerous papers on the solar photospheric abundance of iron have recently been published leading to a longstanding debate concerning rather different results obtained from the analyses of Fe I lines and, to a lesser extent, of Fe II lines. Based on a set of 65 solar Fe I lines, with accurate transition probabilities as well as new accurate damping constants, we construct a new empirical photospheric model. We succeed to reconcile abundance results obtained from low and high excitation Fe I lines as well as from Fe II lines and derive a solar photospheric abundance of iron, $A_{Fe} = 7.50 \pm 0.05$, which perfectly agrees with the meteoritic value.

Key words: atomic data – Sun: abundances – Sun: photosphere

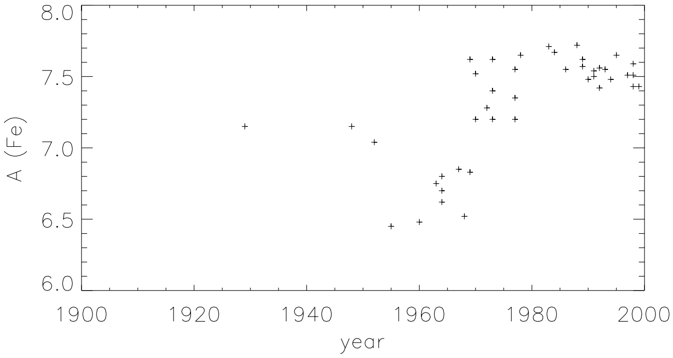

Fig. 1. The iron solar abundance as a function of time, since the pioneering determination by H.N. Russell in 1929. The value obtained by Russell was not too far away from the most recent results of the last decade!

1. Introduction

Although Fe II ions are about 10 times more abundant than Fe I atoms in the solar photosphere, the solar photospheric spectrum does show more than 10 times more Fe I lines than Fe II lines. More trust is generally placed in abundance derived from the dominant stage of ionization because the results hardly depend on temperature. Suitable solar Fe II lines exist, but accurate transition probabilities are available only for very few of these lines (see Sect. 4). Therefore Fe I lines, in a rather large range of excitation energies and with accurate transition probabilities, have commonly been used to derive the solar abundance of iron.

Early determinations resulted in values an order of magnitude less than the coronal and meteoritic abundances. This led to a large number of new mechanisms for producing such a large fractionation in the solar nebula and in the corona. It was finally discovered in 1969 that the transition probabilities of Fe I lines used in these determinations were erroneous i.e. too large by one order of magnitude! The history of the variations of the solar abundance of iron since its first measurement by Russell (1929) is illustrated in Fig. 1.

Recent "oscillations" of the photospheric abundance of iron are discussed in different recent papers devoted to the analysis of Fe I lines (Blackwell et al. 1984; Holweger et al. 1991; Milford et al. 1994; Blackwell et al. 1995a; Holweger et al. 1995; Blackwell et al. 1995b; Kostik et al. 1996; Anstee et al. 1997) as well as Fe II lines (Pauls et al. 1990; Holweger et al. 1990; Biémont et al. 1991; Hannaford et al. 1992; Raassen & Uylings 1998a; Schnabel et al. 1999) or both (Lambert et al. 1996).

Results obtained from Fe II lines generally agree with the meteoritic abundance (see however Sect. 4) but with uncomfortably large uncertainties approaching 25%!

Results obtained from Fe I lines using the same photospheric model (Holweger & Müller 1974) led to a debate between the Oxford group (D.E. Blackwell and co-workers) and the Kiel-Hannover group (H. Holweger, M. Kock and co-workers) as to whether the solar abundance[1] of iron is high, A_{Fe} = 7.63 (Oxford), i.e. larger than the meteoritic value A_{Fe} = 7.50 (Anders & Grevesse 1989; Grevesse & Sauval 1998), or low (Kiel-Hannover), i.e. in agreement with the meteorites. We have to point out here that the Oxford group uses rather low excitation Fe I lines for which the gf-values have been accurately measured at Oxford whereas the Kiel-Hannover group uses Fe I lines of higher excitation which have been accurately measured at Hannover.

The reasons for this longstanding puzzling difference are to be found (see the hereabove mentioned most recent Fe I pa-

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[*] A detailed version of Table 2 is available in electronic form at the CDS via anonymous ftp to cdsarc.u-strasbg.fr (130.79.128.5) or at the ORB via anonymous ftp to ftpserver.oma.be/pub/astro/jacques.

[1] Solar abundances are given in the logarithmic scale usually adopted by astronomers, $A_{el} = \log N_{el}/N_H + 12$, where N_{el} is the abundance by number.

pers) in cumulative effects on the abundance results of slight differences between the equivalent widths, the gf-values absolute scales, the microturbulent velocities and the empirical enhancement factors of the damping constants adopted by the two groups. These effects are small individually, but they are all of the same sign and together they amount to the difference between the "high" and "low" solar iron abundance. But this does not tell us who is right!

Recently, Anstee & O'Mara (1995), Barklem & O'Mara (1997) and Barklem et al. (1998a) (see also Barklem et al. 1998b) have made a decisive progress. They have computed accurate cross-sections for the broadening of s–p, p–s, p–d, d–p, d–f and f–d transitions of neutral atoms by collisions with neutral hydrogen. This allows for the first time to get rid of one of the most uncertain parameters, the so-called enhancement factor, applied for decades to the conventionnally used way to compute the collisional damping constant. Anstee et al. (1997) successfully applied their new cross-sections to the analysis of the wings of very strong solar Fe I lines, never used before for abundance diagnostics. This leads to a very accurate abundance in perfect agreement with the meteorites.

As the broadening parameter plays a non negligible role in the difference Oxford versus Kiel-Hannover, we decided to make a new analysis of a sample of Fe I lines including non blended lines for which accurate transition probabilities have been measured either at Oxford for low excitation lines or at Hannover for higher excitation lines and for which accurate damping constants can be computed from the hereabove mentioned new theory. The results are presented in Sect. 2. Unfortunately, the new damping constants do not help solving entirely the iron problem. In Sect. 3 we propose an alternative solution to this problem with a slightly modified photospheric model which we applied successfully to Fe II lines in Sect. 4. Conclusions are presented in Sect. 5.

2. New analysis of solar Fe I lines

We selected a sample of 65 good solar Fe I lines. The solar data (W_λ, equivalent widths) as well as the atomic data (excitation potentials, log gf-values) are given in Table 1. Among these 65 lines, 33 lines of rather low excitation energy (0–2.6 eV) have gf-values measured at Oxford (see Blackwell et al. 1995a, for the references) whereas for 32 lines of higher excitation energy (1.6–4.6 eV), the gf-values have been measured at Hannover (Bard et al. 1991; Bard & Kock 1994). Broadening of these lines is essentially due to collisions with atomic hydrogen atoms. Damping constants have been computed from the recent data cited above (Anstee & O'Mara 1995; Barklem & O'Mara 1997; Barklem et al. 1998a).

The problem of the accuracy of the two sets of gf-values has been discussed at large in the original papers and also in Blackwell et al. (1995a, b) and Holweger et al. (1995). For the Oxford data, we shall adopt a conservative estimate for the mean uncertainty of the oscillator strengths of Δlog gf (Oxford) = 0.02 dex (= 5%); in reality the uncertainty is probably better than this value (see Blackwell et al. 1995b). The gf-values measured at Hannover have generally uncertainties of the order of 10%. Actually, the mean uncertainty of the log gf-values for our 32 Hannover Fe I lines is Δlog gf (Hannover) = 0.045 (± 0.015) dex. Oxford and Hannover data sets have lines in common: a comparison shows that Δlog gf (Hannover – Oxford) = +0.026 ± 0.044 dex (Blackwell et al. 1995b). This excellent agreement between the two sets, obtained by entirely different techniques, is a pretty good test of the quality of the data in both sets (see however Sects. 3 and 4).

We have to comment on two other recent sets of transition probabilities for Fe I. O'Brian et al. (1991) have measured a large number of gf-values for Fe I lines of different excitation energies. We do not select these lines because the uncertainty of individual results is too large. The comparison of the Hannover and O'Brian et al. results shows that Δlog gf (Hannover – O'Brian) = +0.003 ± 0.15 (Bard & Kock 1994). If the mean absolute scales agree, the spread is uncomfortably large! We also showed that if a large number of good solar lines with gf-values from O'Brian et al. are used to derive the solar abundance of iron, the spread of the results is extremely large (see Fig. 4 of Grevesse & Noels 1993 and Grevesse et al. 1995) and essentially due to uncertainties in the gf-values. Milford et al. (1989) also measured intensities of a sample of Fe I lines relative to reference lines for which the absolute gf-values have been measured by others. Some of the reference lines have uncertain accuracy and moreover some of the reference lines are too far away from the line of interest. Although some of their new results are very accurate, essentially those referred to lines measured at Oxford, other results are of lower accuracy. The use of some of their lines of rather high excitation energy by Milford et al. (1994) to derive the solar iron abundance leads to $A_{\rm Fe}$ = 7.54 ± 0.05, a result in agreement with the meteoritic value. No line from this list is used in the present study.

A comment has also to be made on the equivalent widths because differences in W_λ for lines in common between Oxford and Kiel-Hannover can explain part of the difference between the results of the two groups. The problem is already discussed at large by Blackwell et al. (1995a, b), Holweger et al. (1995) and Kostik et al. (1996). Our equivalent widths, measured on the Jungfraujoch atlas of the solar photospheric spectrum of Delbouille et al. (1973), agree pretty well with the W_λ's obtained at Oxford [W_λ(this work)/W_λ(Oxford) = 0.990 ± 0.015] as well as with the W_λ's measured at Kiev by Kostik et al. [W_λ(this work)/W_λ(Kiev) = 1.002 ± 0.026]. When the same comparison is made with the equivalent widths from Kiel, we find W_λ(this work)/W_λ(Kiel) = 1.06 ± 0.07 (for faint lines, the value is 1.08 ± 0.08). Such a large disagreement has much larger effects on the abundance: for medium-strong lines, the difference might amount to 0.10 dex (25%). We feel confident in our measurements because of the very good agreement with Oxford and Kiev and the low dispersion of the results and because in a few cases, it is evident that the Kiel equivalent widths are too small. But, as the Kiel measurements are not made on the same solar photospheric spectrum as the others, maybe the sun is really a slightly variable star.

Table 1. Fe I lines in the solar spectrum

λ (nm)	log gf	Excit. (eV)	W_λ (mÅ)	$A^3_{\rm Fe}$	Weight	λ (nm)	log gf	Excit. (eV)	W_λ (mÅ)	$A^3_{\rm Fe}$	Weight
438.9251	-4.583^2	0.0516	71.2	7.482	2	632.2694	-2.426^2	2.5881	79.2	7.587	3
444.5476	-5.441^2	0.0873	38.5	7.529	2	635.3840	-6.477^2	0.9100	1.4	7.521	1
522.5533	-4.789^2	0.1101	70.3	7.478	2	648.1878	-2.984^2	2.2786	64.2	7.534	1
524.7059	-4.946^2	0.0873	65.6	7.494	2	649.8946	-4.699^2	0.9582	43.7	7.526	3
525.0217	-4.938^2	0.1213	65.5	7.518	1	651.8374	-2.450^1	2.8300	57.0	7.410	2
532.6145	-2.071^1	3.5700	34.5	7.426	1	657.4233	-5.004^2	0.9900	25.7	7.498	2
541.2788	-1.716^1	4.4400	19.1	7.486	2	658.1214	-4.680^1	1.4800	17.0	7.441	1
549.1835	-2.288^1	4.1900	12.6	7.512	3	659.3880	-2.422^2	2.4327	86.5	7.532	3
560.0227	-1.420^1	4.2600	36.5	7.444	2	660.9119	-2.692^2	2.5592	65.3	7.526	3
566.1348	-1.756^1	4.2800	22.0	7.457	3	662.5027	-5.336^2	1.0100	13.5	7.500	2
569.6093	-1.720^1	4.5500	13.4	7.390	2	666.7721	-2.112^1	4.5800	8.9	7.565	3
570.1553	-2.216^2	2.5600	86.0	7.568	2	669.9142	-2.101^1	4.5900	8.0	7.511	3
570.5468	-1.355^1	4.3000	39.7	7.477	3	673.9524	-4.790^1	1.5600	10.9	7.396	2
577.8458	-3.440^2	2.5900	20.2	7.457	3	675.0160	-2.621^2	2.4242	76.2	7.523	3
578.4661	-2.530^1	3.4000	25.4	7.455	2	675.2711	-1.204^1	4.6400	37.3	7.531	3
585.5080	-1.478^1	4.6100	21.7	7.453	2	679.3266	-2.326^1	4.0800	12.6	7.470	3
590.9978	-2.587^1	3.2100	30.0	7.424	1	680.4003	-1.496^1	4.6500	21.4	7.466	2
595.6700	-4.605^2	0.8590	50.3	7.511	3	680.4277	-1.813^1	4.5800	14.9	7.520	2
608.2715	-3.573^2	2.2227	33.8	7.525	2	683.7009	-1.687^1	4.5900	17.7	7.495	3
612.0255	-5.970^2	0.9100	4.8	7.577	2	685.4828	-1.926^1	4.5900	12.7	7.557	2
613.7000	-2.950^2	2.1979	63.7	7.457	2	694.5214	-2.482^2	2.4242	83.2	7.492	3
615.1622	-3.299^2	2.1759	48.5	7.485	3	697.1936	-3.340^1	3.0200	12.5	7.445	1
617.3343	-2.880^2	2.2227	68.1	7.499	1	697.8861	-2.500^2	2.4844	79.9	7.501	3
618.0208	-2.586^1	2.7300	56.0	7.460	1	711.2173	-2.990^1	2.9900	33.0	7.600	1
620.0320	-2.437^2	2.6085	75.4	7.559	3	718.9156	-2.771^1	3.0700	37.5	7.542	1
621.9289	-2.433^2	2.1979	91.0	7.474	1	740.1689	-1.599^1	4.1900	41.9	7.548	3
623.2649	-1.223^1	3.6500	88.4	7.432	2	772.3214	-3.617^2	2.2800	38.5	7.589	3
624.0652	-3.233^1	2.2200	47.6	7.442	2	791.2871	-4.848^2	0.8590	46.8	7.502	2
626.5141	-2.550^2	2.1759	86.7	7.479	3	807.5158	-5.062^2	0.9100	33.1	7.514	2
627.1283	-2.703^1	3.3300	20.5	7.406	2	820.4107	-6.052^2	0.9100	5.4	7.573	1
628.0624	-4.387^2	0.8590	61.8	7.478	2	829.3522	-2.175^1	3.3000	58.3	7.451	3
629.7801	-2.740^2	2.2227	74.8	7.474	3	836.5644	-2.037^1	3.2500	72.0	7.485	2
631.1505	-3.141^1	2.8300	27.0	7.525	1						

1. log gf-values measured by the Kiel-Hannover group (Bard et al. 1991; Bard & Kock 1994)
2. log gf-values measured by the Oxford group (see Blackwell et al. 1995a for the references)
3. Present work adopting the new photospheric model and ξ = 0.8 km s^{-1} (see text).

We have redone the abundance determination using the Fe I *lines of Table 1 and the widely used, since many decades, photospheric model of Holweger & Müller* (1974). *The results are shown in Fig. 2 where a strong dependence against the excitation energy is clearly visible: lines with low excitations lead to results substantially higher than the high excitation lines which lead to an abundance in agreement with that of the meteorites* ($A_{\rm Fe}$ = 7.50).

The last but very important parameter in abundance studies, namely the microturbulent velocity, has been varied between 0.85 and 1 km s^{-1}. In going from ξ = 0.85 km s^{-1}, the value preferred by the Oxford group, to 1 km s^{-1}, the value favoured by the Kiel-Hannover group, the abundance might decrease by 0.08 dex for medium-strong lines. In the plot of Fig. 2, the detailed results are given for ξ = 0.85 km s^{-1}, but the excitation dependence is not drastically changed if ξ = 1 km s^{-1} is used. It is true however that the lines of the Oxford group are more sensitive to the microturbulent velocity than the lines from the Kiel-Hannover group: $\Delta A_{\rm Fe}$(Oxford) = -0.040 (7.617–7.577) whereas $\Delta A_{\rm Fe}$(Kiel-Hannover) = -0.012 (7.513–7.501) if ξ is increased from 0.85 to 1 km s^{-1}.

3. New photospheric model

We thus face a problem. *Using the best data we could adopt for the equivalent widths, the gf-values, the damping constants, the microturbulent velocity and the model atmosphere, the iron abundance, derived from Fe* I *lines of different excitation energies* (0 to 4.55 eV), *strongly depends on the excitation energy* (see Fig. 2). This dependence cannot be explained by possible small non-LTE effects (Steenbock 1985; Holweger 1988, 1996).

We know that the low excitation lines are generally formed higher in the photosphere than high excitation lines. Fig. 3 is the

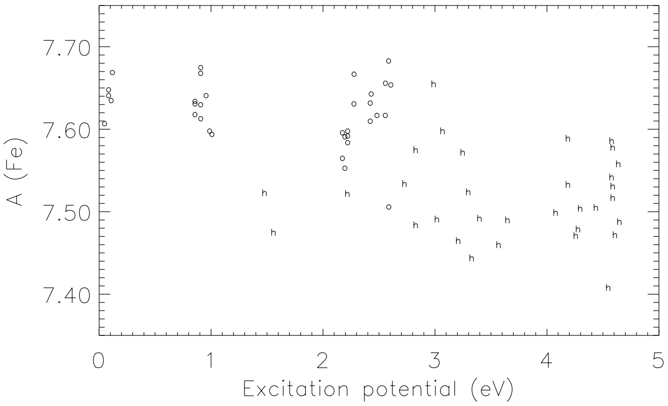

Fig. 2. The iron solar abundance values against the excitation potential for our 65 solar Fe I lines as derived with the Holweger & Müller model (1974) using ξ = 0.85 km s^{-1}. Representative points are indicated ‘o’ for the Oxford set of lines and ‘h’ for the Kiel-Hannover set.

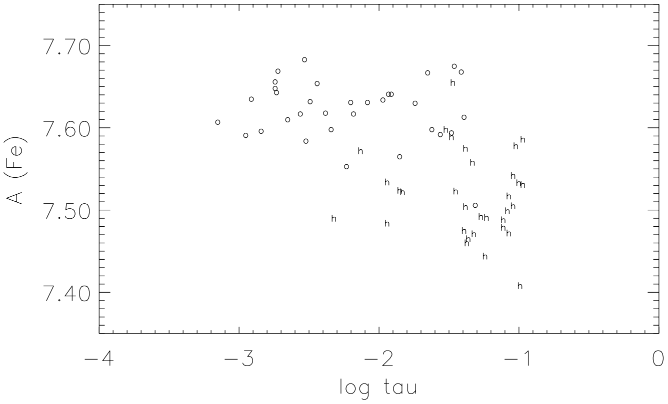

Fig. 3. The iron solar abundance values against the mean optical depth of line formation (log τ_o^*) for our 65 solar Fe I lines as derived with the Holweger & Müller model (1974) using ξ = 0.85 km s^{-1}. Representative points are indicated ‘o’ for the Oxford set of lines and ‘h’ for the Kiel-Hannover set.

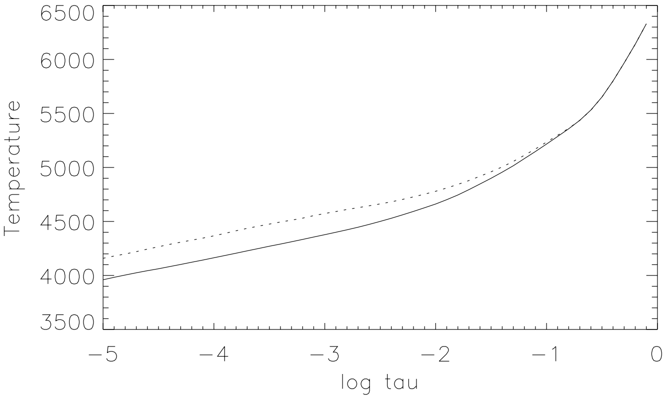

Fig. 4. Temperature distribution against log τ_o for the Holweger & Müller model (1974) (dotted line) and the new photospheric model (solid line).

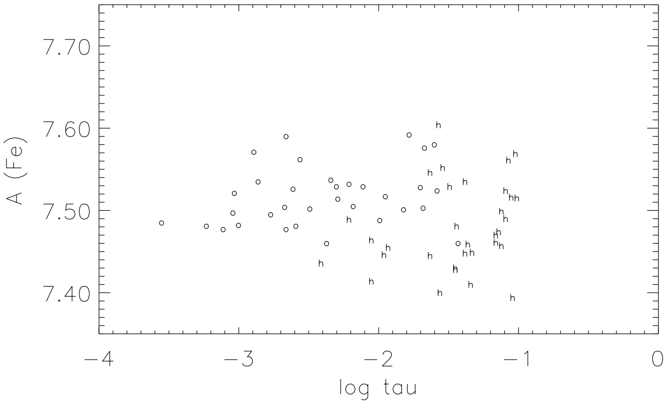

Fig. 5. The iron solar abundance values against the mean optical depth of line formation (τ_o^*) for our 65 solar Fe I lines as derived with our new model (see text). Representative points are indicated ‘o’ for the Oxford set of lines and ‘h’ for the Kiel-Hannover set.

same plot as Fig. 2 but as a function of the mean optical depth of formation of the line (τ_o^*) defined in a way suggested by Magain (1986). We see that most of the Oxford low excitation lines are formed higher in the photosphere than the Kiel-Hannover high excitation lines. As we know that low excitation lines are much more sensitive to the temperature than high excitation lines, *it is tempting to solve the dependence of* $A_{\rm Fe}$ *on excitation and on log* τ_o^* *by slightly decreasing the temperature of the Holweger* & *Müller* (1974) *model in the layers of concern* in order to bring the results obtained with the low excitation lines down to the level of the high excitation lines without changing the result obtained from these lines. Note that we already suggested such a solution to the so-called solar iron problem (Grevesse et al. 1995; Grevesse & Sauval 1998).

We finally came up with the new model illustrated in Fig. 4 and listed in Table 2 (the full table is available at the CDS, but part of the table is included in the paper). Our new model started from the original Holweger & Müller (1974) for layers deeper than log $\tau_o \cong$ -1.0. For higher layers we had to decrease the temperature gradually to reach a uniform ΔT = -200 K for layers higher than log $\tau_o \cong$ -2.0. *With this new model, the solar iron abundance derived from the Fe I lines of Table 1 does not depend any more neither on the mean optical depth of formation nor on the excitation energy* as illustrated by Figs 5 and 6. We also see, when comparing Fig. 6 and Fig. 2, that the results from high excitation lines, hardly depend on the temperature modification.

We derived the best microturbulent velocity in a way suggested by Blackwell et al. (1984): the best value being chosen as the value which leads to the smallest dispersion of the abundance results. We found ξ = 0.8 km s^{-1}, very slightly smaller than Blackwell et al. (1984). With this microturbulent velocity and our new photospheric model, we find a solar abundance of iron of $A_{\rm Fe}$ = 7.497 ± 0.048. We also note that the abundance anomaly for the 2.2 eV lines, found by Blackwell et al. (1984), has disappeared.

The result from the 33 Oxford lines, $A_{\rm Fe}$ = 7.514 ± 0.036, is 0.035 dex larger than the result from the 32 Kiel-Hannover lines, $A_{\rm Fe}$ = 7.479 ± 0.050. This small difference might possibly be related to the 0.026 (± 0.044) dex differences in the ab-

Table 2. New photospheric model

log τ_o	T	log P_e	log P_g
-5.00	3960	-1.767	2.324
-4.50	4062	-1.455	2.641
-4.00	4164	-1.159	2.942
-3.50	4271	-0.872	3.232
-3.00	4378	-0.591	3.517
-2.50	4501	-0.309	3.799
-2.00	4662	-0.017	4.078
-1.50	4900	0.298	4.357
-1.00	5217	0.640	4.634
-0.50	5652	1.062	4.899
0.00	6529	1.850	5.087
0.50	7670	2.731	5.190
1.00	8721	3.361	5.254
1.50	9622	3.803	5.317

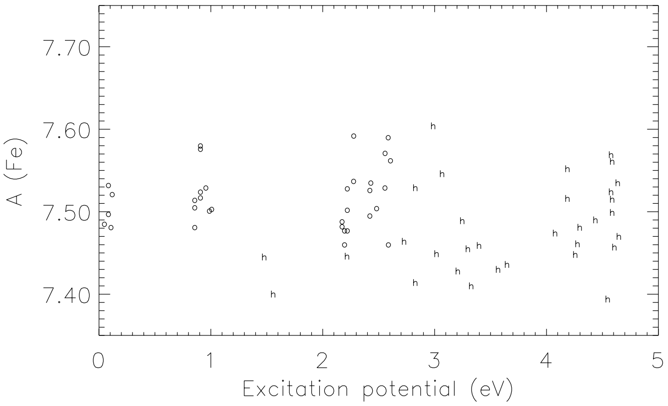

Fig. 6. The iron solar abundance values against the excitation potential for our 65 solar Fe I lines as derived with our new model (see text). Representative points are indicated 'o' for the Oxford set of lines and 'h' for the Kiel-Hannover set.

solute scales of the oscillator strengths between the two groups (Sect. 2).

As we mention hereabove, our new model results from small modifications of the Holweger & Müller (1974) model; moreover this slight temperature decrease occurs in rather high photospheric layers. This new model will not lead to appreciable modifications in the absolute intensities predicted at the center of the solar disk. It will also not change appreciably the center to limb predictions. In other words, as the Holweger & Müller model satisfies these two criteria, i.e. agreement between predicted and observed absolute intensities and center to limb variations respectively, our new model also does.

We have to note that our new model, slightly cooler than the Holweger & Müller model, agrees, at least qualitatively, with a new semiempirical solar photospheric model constructed by Allende Prieto et al. (1998) from the direct inversion of observed line profiles.

4. Fe II lines and other indicators

4.1. *Fe* II *lines*

Results from Fe II lines (Holweger et al. 1990; Biémont et al. 1991; Hannaford et al. 1992) are in good agreement with the meteoritic value. However, other recent analyses (Pauls et al. 1990; Raassen & Uylings 1998a; Schnabel et al. 1999) cast some doubt on this agreement. The first two works conclude to "high" values for the iron abundance: $A_{\rm Fe}$ = 7.66 ± 0.06 and $A_{\rm Fe}$ = 7.59 ± 0.06 respectively, whereas the last authors derive a "low" solar iron abundance: $A_{\rm Fe}$ = 7.42 ± 0.09, even lower than the meteoritic value! Are we going to face for Fe II lines, the same situation as for Fe I lines? The answer is no.

The work of Pauls et al. (1990) has been revised by Hannaford et al. (1992) and reasons given for the "high" result.

The analysis of Raassen & Uylings (1998a) is based on new theoretical semi-empirical transition probabilities calculated using the orthogonal operator approach. These authors show that their theoretical results agree well with recently measured oscillator strengths and lifetimes (see also Raassen & Uylings 1998b). But the tests only concern strong lines in the ultraviolet as well as lifetimes, which are essentially determined by the stronger lines in the branches. The solar lines of interest for abundance analyses are one to two orders of magnitude fainter than the lines tested. When the theoretical oscillator strengths of the solar lines, with low log gf- values around -3, are tested against the best gf-values obtained for Fe II lines using lifetimes and branching fraction measurements (see e.g. Hannaford et al. 1992; Schnabel et al. 1999), it appears that the mean differences, from 11 lines in common, is large: Δlog gf (experiment – theory) = +0.11 ± 0.06. We thus suspect that, even if the theoretical gf-values for strong lines are correct, the relevant data for much fainter solar lines are too small by a rather large amount explaining the "high" abundance value found by Raassen & Uylings (1998a).

The same type of comments has also been made in the past concerning the results of Biémont et al. (1991) who also used solar Fe II lines with semi empirical theoretical gf-values of Kurucz (1988) to derive the solar abundance of iron (Hannaford et al. 1992; Grevesse & Noels 1993; Grevesse et al. 1995). Theoretical techniques have unfortunately not yet been able to achieve the accuracies of experimental techniques for faint lines of heavy elements.

Schnabel et al. (1999) rescaled the experimental gf-value of Heise & Kock (1990) with new very accurate lifetime data and also rescaled the solar abundance analysis of Holweger et al. (1990) to these new gf-values. The new $A_{\rm Fe}$, 7.42 ± 0.09, is somewhat puzzling.

We made a new analysis of Fe II lines with the lines shown in Table 3. Actually the choice of solar Fe II lines is limited by the very few lines for which the oscillator strength is known with accuracy. We added two lines from Hannaford et al. (1992) at 722.2 and 722.4 nm because their gf-values perfectly agree with those of Schnabel et al. (1999) and deleted two lines difficult to measure in the solar spectrum from Schnabel et al.'s list (562.7 and 744.9 nm). For the 13 lines in Table 3, our equivalent widths agree with those of Holweger et al. (1990). With our new model

Table 3. Fe II lines in the solar spectrum

λ (nm)	log gf^1	Excit. (eV)	W_λ (mÅ)	A^2_{Fe}	Weight
457.6339	−2.90	2.844	66.0	7.460	1
462.0520	−3.19	2.828	56.0	7.496	2
465.6981	−3.57	2.891	34.0	7.397	2
523.4630	−2.22	3.221	87.0	7.464	2
526.4790	−3.23	3.250	47.2	7.645	3
541.4075	−3.48	3.221	27.5	7.405	2
552.5130	−3.94	3.267	13.0	7.445	2
643.2683	−3.51	2.891	43.0	7.444	3
651.6083	−3.38	2.891	57.0	7.608	3
722.2397	−3.36	3.890	20.5	7.632	2
722.4464	−3.28	3.890	21.0	7.567	1
751.5837	−3.37	3.903	15.0	7.457	2
771.1731	−2.45	3.903	51.0	7.432	3

1. log gf-values obtained by Schnabel et al. (1999); values for the lines 722.2 and 722.4 nm are from Hannaford et al. (1992)
2. Present work adopting the new photospheric model, ξ = 0.8 km s^{-1}, and an enhancement factor of 2 for the collisional damping constant.

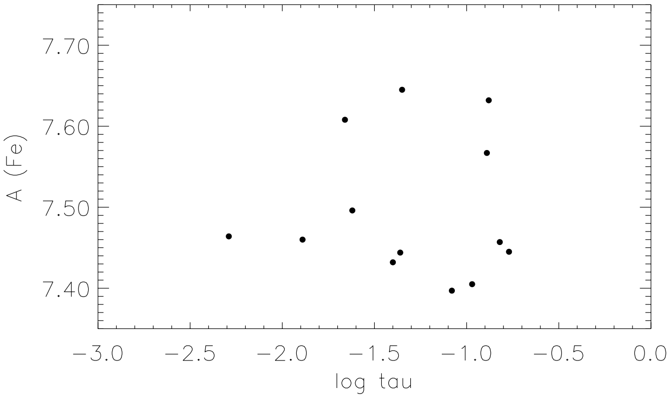

Fig. 7. The iron solar abundance values against the mean optical depth of line formation (τ^*_o) for our 13 solar Fe II lines as derived with our new model (see text).

atmosphere, constructed from the analysis of Fe I lines (Sect. 3), the solar abundance of iron derived from Fe II lines now leads to A_{Fe} = 7.50 ± 0.095, in agreement with the result obtained from the Fe I lines and from the meteorites. However, *the dispersion of the results* (Fig. 7) *is uncomfortably large*.

A few comments have to be made concerning the Fe II lines. The new theory for computing accurate cross-sections for the broadening of spectral lines by collisions with hydrogen atoms (Anstee & O'Mara 1995; Barklem & O'Mara 1997; Barklem et al. 1998a; Barklem et al. 1998b) is not applicable to ions. We therefore used the classical enhancement factor and used different values, 1.5, 2 and 2.5 respectively. Hopefully the lines of Table 3 are not very sensitive to this parameter: when E is enhanced from 1.5 to 2.5, the A_{Fe} decreases by 0.025 dex only. We adopted E = 2 in the results shown in Table 3.

The Fe II results of Table 3 also depend on the microturbulent velocity. When it is decreased from 1 km s^{-1} (Schnabel et al. 1999) to the value we adopted from the Fe I lines, 0.8 km s^{-1}, A_{Fe} increases by 0.03 dex.

Actually, we have thus been able to increase the result of Schnabel et al., A_{Fe} = 7.42 up to 7.50, in agreement with the meteorites, just by making a slightly different choice of lines ($\Delta A_{Fe} \cong +0.02$; the line of Schnabel et al. at 744.9 nm leads to a very small value, 7.25), microturbulent velocity ($\Delta A_{Fe} \cong +0.03$), enhancement factor ($\Delta A_{Fe} \cong +0.01$), the new model atmosphere playing, as expected, a negligible role ($\Delta A_{Fe} \cong +0.01$).

The very large uncertainty of the Fe II *results is certainly related to the mean uncertainty of the oscillator strengths*. This is something which is not new. This is clearly shown by the Fe I and Fe II results when comparing the mean uncertainty of the log gf-values of the lines and the dispersion of the abundance results. In the case of Fe I, for the Oxford lines, Δlog gf = 0.02 dex and ΔA_{Fe} = 0.037 dex whereas for the Kiel-Hannover lines, Δlog gf = 0.045 dex and ΔA_{Fe} = 0.050 dex. For Fe II, the best oscillator strengths available are unfortunately less accurate with Δlog gf = 0.066 dex leading to ΔA_{Fe} = 0.095! A detailed discussion of the oscillator strengths available for Fe I and Fe II lines has recently been made by Lambert et al. (1996).

4.2. Other indicators

The solar photospheric abundance of iron has also been derived from other indicators namely the forbidden lines of Fe II, very faint and difficult to measure with some accuracy in the solar spectrum (Grevesse & Swings 1969), the very high excitation lines of the 4f–5g, 4f–6g and 5g–6h transitions of Fe I in the far infrared (Johansson et al. 1994; Schoenfeld et al. 1995) and also coronal matter which, under certain circumstances, shows photospheric abundances (Feldman 1992).

The present result based on our new model agrees with the abundance results derived from these analyses which agree themselves, with rather large uncertainties in some cases, with the low abundance i.e. the meteoritic value. We emphasize that the forbidden Fe II lines already solved in 1969 the solar iron abundance problem existing at that time (see Sect. 1)!

When our new model is used with the wings of very strong Fe I lines as suggested by Anstee et al. (1997), the abundance result decreases somewhat below the meteoritic value. This is not surprising and probably related to the difficulties one has to reproduce line profiles with homogeneous models.

5. Conclusions

We hope *we have made a step forward in the 1D-modelling of the photosphere. Our new photospheric model, slightly cooler than the widely used, for almost three decades, Holweger &* Müller (1974) *model, does reconcile results based on low and high excitation Fe* I *lines without changing the result obtained from Fe* II *lines. It allows to come to an end with the debate concerning the high versus low solar abundance of iron: the solar photospheric abundance of iron is low,* A_{Fe} = 7.50 ± 0.05; *it agrees with the very accurate meteoritic value*.

The next step has however to abandon the 1D hypothesis in order to describe more realistically the photospheric structure. Actually, we do believe that the Holweger & Müller type photospheric models might possibly be slightly different if one wants to reproduce the behaviour of different tracers of the temperature, like infrared CO lines rather than Fe I lines. *Each indicator of the temperature in the outer photospheric layers might lead to a slightly different homogeneous model* because these different indicators have their own sensitivities to inhomogeneities and will react in a different way to the actual heterogeneous structure of the photospheric layers.

We know that the photospheric layers, just above the convection zone, are very heterogeneous and that matter motions extend very high through the photosphere up to about 1000 km above the top of the convection zone. Line shifts and asymmetries of profiles of unblended lines resulting from these matter motions are observed even on infrared CO lines (Grevesse & Sauval 1991; Blomme et al. 1994; Grevesse & Sauval 1994) which are formed very high into the so-called COmosphere (Ayres 1998). We do believe that we are now really facing the limitations of one-component model atmospheres and have now to turn to more realistic modelling (see e.g. Solanki 1998; Rutten 1998). But the effects on the solar iron abundance are still difficult to estimate (Holweger 1988, 1996; Holweger et al. 1990; Anstee et al. 1997; Gadun & Pavlenko 1997).

A last general comment has to be made concerning transition probabilities. Iron is by far the dominant species in the solar visible spectrum; this is true as well for a large number of stellar spectra. In each angström of the solar spectrum, we find an iron line. At the end of this Millennium, *it is very disappointing to realize that accurate transition probabilities are known for only a minority of these lines*.

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References

Allende Prieto C., Ruiz Cobo B., Garcia Lopez R.J., 1998, ApJ 502, 951

Anders E., Grevesse, N., 1989, Geochim. Cosmochim. Acta 53, 197

Anstee S.D., O'Mara B.J., 1995, MNRAS 276, 859

Anstee S.D., O'Mara B.J., Ross J.E., 1997, MNRAS 284, 202

Ayres T.R., 1998, In: Deubner F.L., Christensen-Dalsgaard J., Kurtz D. (eds.) Proc. IAU Symp. 185, New Eyes to See Inside the Sun and the Stars. Kluwer, Dordrecht, p. 403

Bard A., Kock M., 1994, A&A 282, 1014

Bard A., Kock A., Kock M., 1991, A&A 248, 315

Barklem P.S., O'Mara B.J., 1997, MNRAS 290, 102

Barklem P.S., O'Mara B.J., Ross J.E., 1998a, MNRAS 296, 1057

Barklem P.S., Anstee S.D., O'Mara B.J., 1998b, Publ. Astron. Soc. Aust. 15, 336

Biémont E., Baudoux M., Kurucz R.L., Ansbacher W., Pinnington E.H., 1991, A&A 249, 539

Blackwell D.E., Booth A.J., Petford A.D., 1984, A&A 132, 236

Blackwell D.E., Lynas-Gray A.E., Smith G., 1995a, A&A 296, 217

Blackwell D.E., Smith G., Lynas-Gray A.E., 1995b, A&A 303, 575

Blomme R., Sauval A.J., Grevesse N., 1994, In: Rabin D.R., Jefferies J.T., Lindsey C. (eds.) Proc. IAU Symp. 154, Infrared Solar Physics. Kluwer, Dordrecht, p. 533

Delbouille L., Roland G., Neven L., 1973, Photometric Atlas of the Solar Spectrum from 3000 to 10000 Å. Institut d'Astrophysique, Université de Liège

Feldman U., 1992, Physica Scripta 46, 202

Gadun A.S., Pavlenko Ya.V., 1997, A&A 324, 281

Grevesse N., Noels A., 1993, Physica Scripta T47, 133

Grevesse N., Sauval A.J., 1991, In: Jaschek C., Andrillat Y. (eds.) The Infrared Spectral Region of Stars. Cambridge University Press, p. 215

Grevesse N., Sauval A.J., 1994, In: Jørgensen U.G. (ed.) Molecules in the Stellar Environment. Proc. IAU Coll. 146, Lecture Notes in Physics 428, Springer-Verlag, p. 196

Grevesse N., Sauval A.J., 1998, In: Frölich C., Huber M.C.E., Solanki S.K., von Steiger R. (eds.) Solar Composition and its Evolution - From Core to Corona. Space Science Series of ISSI, Kluwer Acad. Publ., p. 161 (also Space Sci. Rev. 85, 161, 1998)

Grevesse N., Swings J.P., 1969, A&A 2, 28

Grevesse N., Noels A., Sauval A.J., 1995, In: Sauval A.J., Blomme R., Grevesse N. (eds.) Laboratory and Astronomical High Resolution Spectra. ASP Conference Series Vol. 81, p. 74

Hannaford P., Lowe R.M., Grevesse N., Noels A., 1992, A&A 259, 301

Heise C., Kock M., 1990, A&A 230, 244

Holweger H., 1988, In: Cayrel de Strobel G., Spite M. (eds) The Impact of Very High S/N Spectroscopy on Stellar Physics. Reidel, Dordrecht, p. 411

Holweger H., 1996, Physica Scripta T65, 151

Holweger H., Müller E.A., 1974, Solar Physics 39, 19

Holweger H., Heise C., Kock M., 1990, A&A 232, 510

Holweger H., Bard A., Kock A., Kock M., 1991, A&A 249, 545

Holweger H., Kock M., Bard A., 1995, A&A 296, 233

Johansson S., Nave G., Geller M., et al., 1994, ApJ 429, 419

Kostik R.I., Shchukina N.G., Rutten R.J., 1996, A&A 305, 325

Kurucz R.L., 1988, In: McNally M. (ed.) Transactions IAU XXB. Kluwer, Dordrecht, p. 168

Lambert D.L., Heath J.E., Lemke M., Drake J., 1996, ApJS 103, 183

Magain P., 1986, A&A 163, 135

Milford P.N., O'Mara B.J., Ross J.E., 1989, JQSRT 41, 433

Milford P.N., O'Mara B.J., Ross J.E., 1994, A&A 292, 276

O'Brian T.R., Wickliffe M.E., Lawler J.E., Whaling W., Brault J.W., 1991, JOSA B8, 1185

Pauls U., Grevesse N., Huber M., 1990, A&A 231, 536

Raassen A.J.J., Uylings P.H.M., 1998a, A&A 340, 300

Raassen A.J.J., Uylings P.H.M., 1998b, J. Phys.B: At. Mol. Opt. Phys. 31, 3137

Russell H.N., 1929, ApJ 70, 11

Rutten R.J., 1998, In: Frölich C., Huber M.C.E., Solanki S.K., von Steiger R.(eds.) Solar Composition and its Evolution - From Core to Corona. Space Science Series of ISSI, Kluwer Acad. Publ., p. 269 (also Space Sci. Rev. 85, 269, 1998)

Schnabel R., Kock M., Holweger H., 1999, A&A 342, 610

Schoenfeld W.G., Chang E.S., Geller M., et al., 1995, A&A 301, 593

Solanki S.K., 1998, In: Frölich C., Huber M.C.E., Solanki S.K., von Steiger R.(eds.) Solar Composition and its Evolution - From Core to Corona. Space Science Series of ISSI, Kluwer Acad. Publ., p. 175 (also Space Sci. Rev. 85, 175, 1998)

Steenbock W., 1985, In: Jaschek M., Keenan P.C. (eds.) Cool Stars with Excesses of Heavy Elements. Reidel, Dordrecht, p. 231