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The Chemical Composition of the Sun

John E. Ross and Lawrence H. Aller

Background

Very extensive spectroscopic and solar wind studies permit some general conclusions about the solar chemical composition and, by inference, the original chemical composition of the solar system. Table 1 gives references (*1–108*) for individual elements. Present data support the suggestion by Goldschmidt (*109*) that chondritic meteorites give the best clues to solar system abundances of nonvolatile elements. The carbonaceous chondrites, which have suffered a minimum of chemical processing, provide the best data. To obtain quantities of volatile elements, however, we must rely on the solar atmosphere.

Although a qualitative analysis of the sun was carried out in the earliest days of spectroscopy, the first quantitative study was undertaken by Russell (*90*), who used eye estimates of the intensities of the dark Fraunhofer lines calibrated by theoretical line strengths. Considering the crudeness of the procedures, his results were remarkably accurate. A great step forward was Minnaert's development (*110*) of the curve of growth technique. The intensity of a line may be expressed in terms of the energy removed from the nearby continuous spectrum, expressed as equivalent angstroms thereof. Essentially, this equivalent width (usually divided by the wavelength) is plotted against the product of the number of atoms capable of absorbing the line multiplied by the Ladenburg *f* or oscillator strength. Early applications were made by Menzel, Unsöld, and others (*111*). Later investigations by Claas (*112*), Weidemann (*113*), and Goldberg *et al.* (*114*) employed model atmospheres in which the explicit variations of temperature and density with depth were treated. These methods, however, handle only part of the information contained in the spectral lines themselves. For weak lines, lines falling in crowded regions of the spectrum, or lines affected by hyperfine structure it is necessary to use a method of spectrum synthesis, such as has been developed by Ross *et al.* (*81*), Hauge and co-workers (*38*), Grevesse (*22*), and others.

Earlier work was confined to investigations of the spectrum of the bright surface or photosphere of the sun. With the advent of rockets and spacecraft, it was possible to use the spectrum of the corona, in the visible, extreme ultraviolet (EUV), and x-ray regions. Pottasch's pioneering investigations (*76*) have been followed by extensive studies by many groups (*14, 17, 71, 105*). Furthermore, it is possible to measure the actual fluxes of particles ejected from the sun in flare activity and "conventional" solar wind.

Basic Methods

Photospheric spectrum. Analysis of the photospheric spectrum yields information for the greatest number of elements, since lines of even some very rare elements such as thorium are found. Here, we emphasize results obtained by the spectrum synthesis method; an account of this and older methods may be found in the literature (*115*).

The basic idea is that an absorption line is formed in the photospheric spectrum because there is a temperature gradient in the solar atmosphere. Spectral lines appear at wavelengths where the absorptivity is enhanced; the line absorption is added to that of the continuum. At wavelengths corresponding to spectral lines we see only the shallowest, coolest layers because the absorptivity is high. In the neighboring continuum we observe radiation that comes, on the average, from deeper, hotter layers, since the absorptivity there is less.

The most obvious manifestation of the temperature gradient in the solar atmosphere is the darkening to the limb. As one looks toward the edge of the solar disk, the line of sight makes a greater angle to the outward normal—that is, one looks more and more obliquely into the solar atmosphere and hence into shallower layers that shine less brightly because they are cooler. In the outermost layers the temperature rise depends on the opacity of the material, which is primarily determined by processes involving hydrogen. Throughout most of the usually observed solar spectrum, the principal source of continuous absorption is the negative hydrogen ion, whose concentration depends on the local temperature and the electron pressure. Most of the electrons are supplied by the relatively easily ionized metals; hence the metal/hydrogen ratio is important even though direct continuous absorption by metallic atoms themselves is not significant.

At a wavelength λ, the intensity of an emergent ray making an angle θ with respect to the outward normal can be shown to be (*115*)

$$I_\lambda(0,\theta) = \int_0^\infty S_\lambda(\tau_\lambda) e^{-\tau_\lambda \sec\theta} \sec\theta \, d\tau_\lambda \qquad (1)$$

where $S\lambda(\tau\lambda)$ is a monotonically increasing function of the depth. In local thermodynamic equilibrium (LTE), where the population of each atomic level is given by the Boltzmann distribution and Saha equation for the local electron pressure and gas kinetic temperature, $S_\lambda(\tau_\lambda)$ is given by the Planckian equation $B_\lambda[T(\tau_\lambda)]$.

Here, τ_λ is the optical depth. In the continuous spectrum, it is defined by

$$d\tau_\lambda{}^c = \kappa_\lambda{}^c \rho dh \qquad (2)$$

where $\kappa_\lambda{}^c$ is the coefficient of continuous absorption at λ calculated per gram of solar material. It is proportional to the abundance of hydrogen. Here ρ is the density and dh the element of linear depth.

In a spectral line

$$d\tau_\lambda{}^L = (l_\lambda + \kappa_\lambda{}^c) \rho dh \qquad (3)$$

where l_λ is the atomic line absorption coefficient calculated per gram of solar material

$$l_\lambda = n_{r,s}\, \alpha(\lambda_0)\, H\left(a, \frac{\lambda - \lambda_0}{\Delta\lambda_D}\right) \qquad (4)$$

John Ross is on the staff of the Department of Physics, University of Queensland, Queensland, Australia, and Lawrence Aller is professor of astronomy at the University of California, Los Angeles 90024.

where $n_{r,s}$ denotes the number of atoms in the level r of the ionization stage s of a particular element E. Only these atoms of element E are capable of absorbing the line in question. The atomic absorption coefficient at λ_0, the center of the line $\alpha(\lambda_0)$, is proportional to the Ladenburg f or oscillator strength and inversely proportional to $\Delta\lambda_D$, the Doppler width of the line, which is determined by the gas kinetic motion and mass motions in the solar atmosphere (*115*). The term in H is the usual line-broadening function; and a is proportional to the damping constant divided by the Doppler width, all expressed in frequency units. [For a detailed account see Aller (*116*, p. 322).]

In addition, certain lines are broadened by hyperfine structure or by isotope shifts, while in magnetically active areas such as sunspots the Zeeman effect must be taken into account. We may write

$$n_{r,s} = \frac{n_{r,s}}{n_r} \frac{n_r}{n_{\mathrm{E}}} \frac{n_{\mathrm{E}}}{n_{\mathrm{H}}} n_{\mathrm{H}} \qquad (5)$$

where n_r is the total number of atoms of element E in ionization stage r, n_{E} is the number of atoms of the element in question, and n_{H} is the number of hydrogen atoms, all reckoned per gram of solar material. In local thermodynamic equilibrium, $n_{r,s}/n_r$ is given by the Boltzmann distribution, and n_r/n_{E} is computed from the Saha ionization equation, all by using the local temperature and electron pressure. In a nonlocal thermodynamic equilibrium (non-LTE) situation we must consider detailed physical processes acting to populate individual atomic levels. We must know the precise radiation field at each point in the atmosphere, and atomic parameters such as absorption and recombination coefficients, collisional cross sections, and so forth. In any event, at each point in the atmosphere, we must know the electron pressure, P_ϵ, the temperature T, the gas pressure P_g, and the continuum opacity; these quantities define the model of the atmosphere. Since they depend on its chemical composition, which is what we are actually trying to determine, we must proceed by an iterative process.

Because hydrogen is overwhelmingly the most abundant element in the sun and similar stars, it contributes most of the atoms and most of the mass to each gram of solar material. The He/H ratio does not affect the continuous absorptivity per gram of material, but it does affect the total weight and hence the relationship between gas pressure and electron pressure. Furthermore, the He/H ratio is of great interest in connection with the history of the solar system.

In the method of spectrum synthesis, one tries to fit observed shapes of spectral lines. The residual intensity—that is, the ratio of the intensity at any point in a line profile to that in the nearby continuum—is obtained from Eq. 1 as

$$r_\lambda(0,\theta) = I_\lambda{}^{\mathrm{L}}(0,\theta)/I_\lambda{}^{\mathrm{c}}(0,0) \qquad (6)$$

where I_λ^{L} and I_λ^{c} are calculated by using $\tau_\lambda^{\mathrm{L}}$ and $\tau_\lambda^{\mathrm{c}}$ as defined by Eqs. 3 and 2, respectively. Since $\kappa_\lambda^{\mathrm{c}}$ depends on the properties of the negative hydrogen ion and I_λ depends on n_{E}, what we derive is the ratio of the abundance of the given element to that of hydrogen. Therefore, all solar abundances are normalized with respect to hydrogen. Usually we take the logarithm of the abundance of hydrogen $\log A(\mathrm{H}) = 12.00$. If we want to normalize with respect to silicon, we have to select a specific abundance value for this element. Here we take the silicon abundance as $\log A(\mathrm{Si}) = 7.65$, essentially the value proposed by Holweger (*44*).

The photospheric abundances will depend on the model atmosphere chosen. Among popular models are those proposed by Holweger (*117*) and by Elste (*118*), the Harvard-Smithsonian Reference Atmosphere (HSRA) (*119*), and a recent revision thereof by Vernazza *et al.* (*120*). The latter model attempts to take into account the breakdown of local thermodynamic equilibrium in the upper photosphere. The model atmosphere selected can affect abundance determinations. With respect to relative abundances of metals, our choice between the above-mentioned models will affect $\log A$ by typically only 0.01 or 0.02, whereas the absolute abundance of a metal such as nickel may vary by as much as $\log A = 0.1$ in going from the Holweger model to HSRA (*61*).

In the uppermost layers of the photosphere, deviations from local thermodynamic equilibrium can be significant. Slight errors in temperature distribution severely affect the calculated populations of oxygen atoms in levels that give rise to the near-infrared lines of this element.

In the ultraviolet, uncertainties in the boundary temperature play an important role. Furthermore, the opacity here comes mostly from overlapping profiles of discrete spectral lines (*120, 121*).

The kinematical structure of the solar atmosphere is important but complicated [see, for example, the discussion by Ulrich (*122*)]. Mass motions of radiating and emitting gases affect spectral line broadening and therefore the observed line profiles. A wrong model for the velocity field produces false estimates of effects of collisional broadening and erroneous abundances. Just below the outermost layer of the photosphere, which is in radiative equilibrium (that is, the transport of energy there is by radiation), lies a thick zone in turbulent convection. Earlier investigations introduced the concept of microturbulence, envisaging the solar atmosphere as containing turbulent elements whose sizes were small compared with the mean free path of a light quantum. At the opposite extreme of macroturbulence, the inverse situation exists. Actually, the true motions are partly orderly and partly chaotic, and questions may be raised as to whether microturbulence, as defined above, really exists.

The solar atmosphere actually involves a regime of rising and falling streams of gas—in some kind of systematic pattern as exhibited by granulation and supergranulation structures (*123*). The observed motions correspond to a scale appropriate to macroturbulence. However, if there are in these streams of moving material mass motions of velocity with gradients of the order of 1 km/sec per $\Delta \log \tau = 1$—that is, $dV/d \log \tau = 1$ km/sec—the practical effects of microturbulence will be duplicated and there will be no need to appeal to an actual small-scale "turbulence." Further, the horizontal and vertical velocities may differ and there can exist temperature differences between rising and falling elements.

The essential atomic parameters are f-values and collisional damping constants, particularly those corresponding to interactions between radiating atoms and neutral hydrogen atoms (van der Waals interactions).

Although the situation has greatly improved in recent years, a determination of accurate f-values remains one of the most troublesome problems in solar abundance determinations. Theoretical procedures are limited to only a small number of transitions in atoms and ions of fairly simple structure, and must be checked by careful experimental work. Empirical f-values have been found from measurements of emission line intensities in calibrated arcs, from absorption in electric furnaces, from anomalous dispersion measurements, from atomic beam experiments, from determinations of lifetimes of excited levels (particularly from beam-foil experiments), and by other methods.

In the solar spectrum, lines are broadened primarily by collisions of the radiating atoms with neutral hydrogen atoms. In the laboratory, collisional effects can be studied with helium or argon used as the broadening agent, but the results must be extrapolated to hydrogen. Thus, it is preferable to use weaker lines that are not influenced noticeably by collisional broadening.

The effects of deviations from local thermodynamic equilibrium are manifest in observed line intensities of many elements and in the centers of strong lines. For example, lines that arise from the upper ^2P term of NaI show pronounced center-limb effects. Usually, at the center of the disk, the effects are minimal, but it is necessary to pay attention to the term structure and to heed center-limb variations. Ultimately, only results obtained from lines that give the same abundance at the center and at the limb can be trusted. Often spectral synthesis permits a good fit at the center of the disk, but there are discrepancies between theory and observation at the limb. In practice, we prefer to adjust the model to reproduce the center-limb variations of most iron lines. Lines of other elements may still be slightly out of step. An example in point is barium, where calculations by K. Sedwick show that systematically too small residual intensities are predicted for some lines (see Fig. 1). For other lines the converse is true.

Chromospheric and coronal abundances. Some elements such as the noble gases have no absorption lines in the observable portions of the dark-lined Fraunhofer spectrum of the photosphere. Other abundant elements such as iron are represented in the spectra of the chromosphere and corona as well as that of the photosphere.

The chromosphere is the buffer zone between the brilliant photosphere of the sun, which has a brightness temperature near 5800°K, and the tenuous outer corona, which has a gas kinetic temperature in the neighborhood of 2,000,000°K. It has an extremely inhomogeneous structure and exhibits severe departures from local thermodynamic equilibrium. The chromosphere is adequately observable only at times of total solar eclipse.

In the visible region of the spectrum, the solar corona is observable only at times of total solar eclipse or with specially devised equipment (coronagraph). It exhibits a continuous spectrum produced by electron scattering, on which are superposed broad forbidden emission lines of highly ionized metals. In the extreme ultraviolet and x-ray regions occur resonance transitions of ions of abundant elements. Measurements of the intensities and spatial distributions of these radiations are made routinely from satellites. Interpretation of the data depends on construction of coronal models, including an evaluation of the radiation field and relevant atomic data such as cross sections for collisional excitation and ionization, recombination coefficients, and transition probabilities. The procedure resembles that employed for gaseous nebulas, except that for the sun one has essentially complete spectral coverage. The line intensities give certain diagnostic parameters for the ambient plasma, and the construction of a coronal model is carried out by iteration. Fortunately, it is possible to observe certain important elements such as iron in all significant stages of ionization. Hence an excessive dependence on extrapolation procedures can be avoided. There remain, however, uncertainties due to structural inhomogeneities on a scale below what can be observed. Withbroe (*105*) gives a summary of data obtained to 1971.

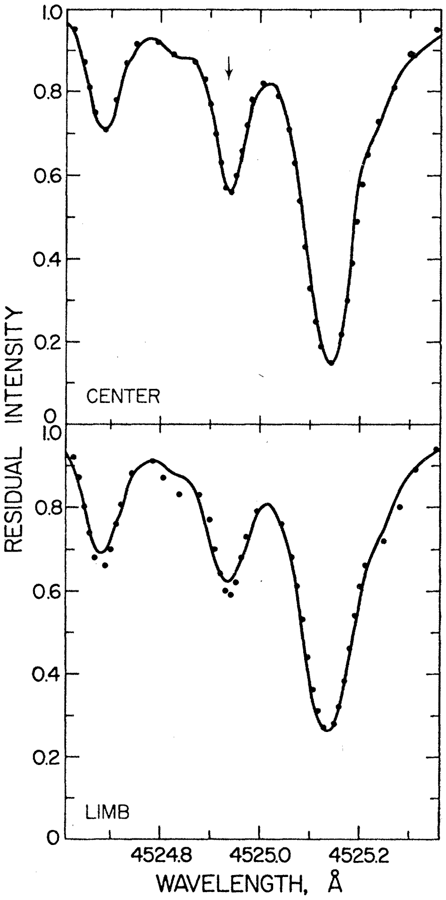

Fig. 1. Application of the spectrum synthesis method to ionized barium, 4524.94 Å. A small portion of the solar spectrum as observed at Kitt Peak National Observatory (solid line) is matched by spectrum synthesis (dots). Once the model atmosphere is adopted, the adjustable parameters are log *fA* (product of *f*-value and abundance) and damping constants for collisional broadening (up to about three times the value calculated by the elementary theory). Slight adjustments to initial wavelengths are permitted. Mean microturbulent and macroturbulent velocities of 1.0 and 2.0 km/sec, respectively, are employed. Data are assessed at the center of the solar disk and near the limb, where $\cos \theta = 1.0$ and 0.2, respectively (θ is the angle between the line of sight and the outward normal). The target line of BaII is indicated by the arrow; other lines include 4524.69 Å of TiII, 4524.84 Å of CrI, and 4525.14 Å of FeI. Additional unidentified lines appear to fall at 4524.74, 4525.02, and 4525.25 Å; these are synthesized as iron lines. Although the FeI line at 4525.14 Å fits well both at the center of the disk and near the limb, thus justifying our choice of model atmosphere, the TiII and BaII lines are predicted to be too deep near the limb; possibly non-LTE effects are important. The CrI feature may be blended. [Courtesy of Katherine Sedwick]

Solar wind, flares, and so forth. Streams of particles are emitted from the sun in the solar wind; high-energy particles, including cosmic rays, are ejected from solar flares. Significant quantitative differences exist from flare to flare. Webber *et al.* (*101*) suggest that not only may the acceleration mechanisms differ from one flare to another, but the number of ions of any particular type ejected may depend on its rigidity. Long-term averages over the solar wind may be more useful. Geiss *et al.* (*124*) have shown that the ratio He/Ne/Ar may be established accurately from the solar wind.

Table of Solar Elemental Abundances

Table 1 lists the elements whose abundances have been determined or estimated in the sun. Several elements are missing. Invariably these elements are either relatively rare or have lines that are masked by transitions of abundant elements such as iron, at least in the visible range of the spectrum. Examples include As, Se, Te, I, and Ta; Kr and Xe have transitions exclusively in the far ultraviolet, where they also appear to be masked.

The first two columns are self-explanatory; they give the atomic number and element. The third column gives the adopted log*A*(element) on the scale log*A*(H) = 12.00, together with the suggested error in the determination. These error estimates represent our best judgment on the basis of existing data; the errors may prove to be larger if there are gross inaccuracies in the *f*-values, as proved to be the situation some years ago for iron. The fourth column gives the quality of the determination. Very few elements would merit an A rating (abundance known to an accuracy of about 10 percent); perhaps oxygen would so qualify. A B represents a reasonably good determination; the limitations here are usually *f*-value uncertainties and non-LTE effects. Typical examples are C, N, Si, K, Ca, Sc, Ti, and perhaps Fe (to which we devote a few remarks below). A C represents a fair determination; some elements such as Ge are rated C even though the formal error is small; because they are represented in the solar spectrum by only a few lines. In other examples, such as Mg, there are grievous discordances between *f*-values for the neutral atom and the ion. A D rating represents a poor determination; the problem may be blends or poor *f*-values. In all instances where the *f*-values are obtained only from calibrated arcs, where both large systematic and accidental errors occur, a rating of D is given unless there is independent supporting evidence.

Column 5 gives the solar system abundance compiled by Cameron (*125*); these data are mostly derived from carbonaceous chondrites, for which we give data in columns 6 and 7 as assembled by Urey (*126*) for both type I and type II carbonaceous chondrites (CI and CII). Note that we take $\log A(\mathrm{Si}) = 7.65$.

Column 8 gives the source of the data as follows.

p) Photospheric absorption lines, which can include lines of neutral atoms or ions, or even forbidden lines, as for [NI], [OI], and [SiI].

ss) Atomic or molecular lines observed in sunspots.

ch) Chromosphere; the data here are taken mostly from Pecker and Pottasch (*75*) and are subject to large uncertainties because of severe deviations from local thermodynamic equilibrium.

c) Corona, which can include lines in the visible, EUV, or x-ray regions (*105*).

w) Solar wind and cosmic rays from flares.

Column 9 gives the source of f-values—T denotes theoretical and E empirical data—while column 10 gives the quality of the f-values. Here A generally denotes values deemed accurate to within 10 percent, B indicates they are probably good to about 15 or 20 percent, C implies a 30 to 50 percent inaccuracy, whereas D quality f-values are subject to large accidental or systematic errors that may amount to factors as high as 4 or 5. In many spectra, the quality of the transition probabilities employed varies over a considerable range.

In column 11 we list notes pertinent to the analysis.

a) Deviations from local thermodynamic equilibrium are or may be important for at least some transitions customarily employed in the analysis. Lines of the heavier rare elements may also be affected by this phenomenon, but "a" is not indicated unless there is independent evidence for it. For very weak lines other effects, particularly blending, are more important.

b) Improved f-values (or in some instances also collisional damping constants) are needed. For relatively few atoms can it be said that the f-value situation is in a truly satisfactory state at the present time.

c) A spectrum synthesis is needed for this element; of course, for some lines this refined procedure is always necessary.

d) Molecular formation is important, either because it depletes the element or because the element appears only in molecular form (for example, Cl).

e) Hyperfine structure of the lines must be considered; under such circumstances a spectrum synthesis must be carried out.

f) The model for the line-forming region in the chromosphere or corona needs improvement; this statement is valid for most coronal lines.

g) Lines are formed in sunspots; hence the results are sensitive to the adopted model for the spots.

h) Line blending is serious.

Underlined letters mean that the state-

Table 1. Solar elemental abundances. Symbols are explained in the text.

Atomic number	Element	Adopted log*A*	Quality of determination	Log*A*			Source of solar data	Source of f-values	Quality of f-values	Notes on analysis	References
				Cameron solar system	Urey						
					CI	CII					
2	He	10.8 ± 0.2	C	11.00			ch, w	T	A	See text	(*18, 41, 52, 73*)
3	Li	1.0 ± 0.1	B	3.35	3.35	3.31	p, ss	T, E	A	a, c, e, h	(*20, 69, 97*)
4	Be	1.15 ± 0.2	B	1.56			p	T, E	A	a, c, h	(*9, 22, 88*)
5	B	<2.1 ± 0.2		4.19	3.78	3.92	p, ss				(*33*)
6	C	8.62 ± 0.12	B	8.72			p, c, w	T	B–C	a, b, d, f	(*13, 17, 25, 53, 66, 70, 71, 74, 84*)
7	N	7.94 ± 0.15	B	8.22			p, c, w	T	B–C	a, b, d, f	(*13, 17, 25, 53, 71, 74, 84, 98, 105*)
8	O	8.84 ± 0.07	A	8.98			p, c, w	T	B–C	a, b, c, d	(*13, 17, 25, 53, 67, 70, 71, 74, 84, 91, 98, 105*)
9	F	4.56 ± 0.33	C	5.04	5.21	5.05	ss	E	C?	b, d, g, h	(*31*)
10	Ne	7.57 ± 0.12	B	8.19			c, w	T		a, f	(*13, 17, 71, 76, 98, 104, 105*)
11	Na	6.28 ± 0.05	A	6.43	6.42	6.20	p, c	T, E	A–B	a, b, f	(*17, 42, 58, 71, 98, 105, 107*)
12	Mg	7.60 ± 0.15	C	7.67	7.67	7.66	p, c, w	T, E	C	a, b	(*13, 17, 42, 60, 71, 74, 98, 105*), R
13	Al	6.52 ± 0.12	C	6.58	6.58	6.57	p, c	T	C–D	a, b, f	(*17, 58, 71, 74, 76, 98 105*)
14	Si	7.65 ± 0.08	B	7.65	7.65	7.65	p, c, w	T, E	C	a, b, f	(*17, 30, 44, 59, 71, 74 76, 105*)
15	P	5.50 ± 0.15	C	5.63	5.75	5.57	p, c	T	C–D	b	(*58, 74, 76, 105*), R
16	S	7.2 ± 0.15	C	7.35	7.35	7.01	p, c, w	T	C	b, c, f	(*13, 17, 60, 71, 74, 76, 96, 105*), R
17	Cl	5.5 ± 0.4	D	5.41	4.95	4.97	ss	T	C	b, c, d, g	(*32*)
18	Ar	6.0 ± 0.20	C	6.72			c, w	T	B–C	a, b, f	(*13, 14, 71, 98*)
19	K	5.16 ± 0.10	B	5.47	5.15	4.97	p, c	T, E	B	a, c, f	(*15, 58, 74, 76*), R
20	Ca	6.35 ± 0.10	B	6.51	6.47	6.51	p, c, w	E	B	a, f	(*13, 14, 43, 58, 63, 71, 74, 105*)
21	Sc	3.04 ± 0.07	B	3.19	3.19	3.17	p	E	B	a?, b	(*6, 25, 39, 100*)
22	Ti	5.05 ± 0.12	B	5.09	5.05	5.03	p, c, w	E	B	a?, b, f	(*5, 13, 25, 68, 75, 106*), R
23	V	4.02 ± 0.15	C	4.07	4.12	4.06	p, c	E	C	a?, b, c, e, f	(*25, 68, 71, 100*)
24	Cr	5.71 ± 0.14	C	5.75	5.73	5.75	p, c	E	B–C	a?, b, f	(*10, 11, 25, 68, 71, 105*), R
25	Mn	5.42 ± 0.16	C	5.62	5.62	5.44	p, c	E	C	a?, b, c, e, f	(*7, 14, 25, 64, 68, 72*)
26	Fe	7.50 ± 0.08	B?	7.57	7.60	7.57	p, c, w	T, E	A–C	a?, b	See text
27	Co	4.90 ± 0.18	C–	5.00	5.01	4.93	p, c	E	C–D	a, b, c, e	(*14, 25, 46, 50, 71, 72*)
28	Ni	6.28 ± 0.09	B	6.33	6.31	6.32	p, c, w	E	A–B	a?, f	(*28, 61, 63, 71, 105*)
29	Cu	4.06 ± 0.13	B	4.38	4.61	4.41	p, c	E	B	a, c, e	(*14, 51*), R
30	Zn	4.45 ± 0.15	C	4.75	4.78	4.45	p	E	B	a?, b, c	(*57*), R
31	Ga	2.80 ± 0.15	C	3.33	3.36	3.10	p	E	B	c, h?	(*57, 83*)

ment is emphasized; for example, e means that hyperfine structure is extremely important (as for rhenium), c means a spectrum synthesis should be carried out, and b means present *f*-values are totally unsatisfactory.

The last column gives the references; if a particular reference is singled out, it is underlined. Various authors have carried out abundance determinations with a variety of atmosphere and *f*-value choices. In a number of instances, with the assistance of C. Keyes, we have recalculated the data with improved *f*-values or with the HSRA model atmosphere (or both) in order to put the determinations in a more uniform system. In such instances the designation R has been added to the references.

We must emphasize that the analyses vary in quality over a large range—from curve of growth measurements interpreted with poor *f*-values (as for those elements denoted by D) to meticulously careful investigations such as Hauge's study (*35*) of Rb or the study of Li by Müller *et al.* (*69*). For many elements, however, further work does not appear justified until better *f*-values are found.

Two Difficult Problems, the Abundances of Helium and Iron

Two elements of crucial astrophysical and cosmochemical importance are helium and iron, for neither of which can our present information be considered fully satisfactory. The resonance lines of helium fall in the far ultraviolet, where we observe only the radiation of the chromosphere and corona. The lines of helium that fall in ordinary spectral regions originate from a lower level, of excitation potential 20.6 ev. Hence they are not observed in the normal photospheric spectrum. These lines may be found in excited, magnetized (plage and sunspot) regions, in flares, and so forth. They are excited in the chromosphere and in solar prominences under extreme non-LTE conditions. Lines of ionized helium appear in the chromosphere-corona interface, but not in the corona itself, where the temperature is too high. Determinations of the solar helium abundance have also been attempted from stellar structure studies, solar cosmic rays, and the solar wind.

Helium abundance estimates from solar prominences require models and elaborate corrections for deviations from local thermodynamic equilibrium. Thus, Hirayama (*41*) found $n(\mathrm{He})/n(\mathrm{H}) = 0.065 \pm 0.015$. Very similar values have been obtained from solar cosmic rays (*18*, *52*), but studies of the solar wind (*73*, *127*) show that the He/H ratio depends on the level of solar activity. Thus, although we can get good "mean solar wind compositions" by averaging over a huge number of events, it is not clear how representative this number is of the outer convective zone of the sun. In other words, fractionation may depend on the acceleration mechanism. It must also be noted that the ratio quoted above, $\simeq 0.06$, is less than the "canonically accepted" value for the He/H ratio in the

Atomic number	Element	Adopted log*A*	Quality of determination	Log*A*			Source of solar data	Source of *f*-values	Quality of *f*-values	Notes on analysis	References
				Cameron solar system	Urey						
					CI	CII					
32	Ge	3.50 ± 0.08	C	3.71	3.78	3.53	p	E, T	B	c	(*87*)
37	Rb	2.60 ± 0.05	B	2.42	2.46	2.26	p	E	A	c, e, h	(*35*, *55*)
38	Sr	2.90 ± 0.10	B	3.08	3.03	3.02	p, ch	E, T	A–B	a?, c	(*36*, *58*, *75*)
39	Y	2.10 ± 0.25	D	2.33	2.31	2.32	p	E	D	a?, b, c	(*12*), R
40	Zr	2.75 ± 0.16	D	3.1	3.06	2.76	p	E	D	a?, b, c	(*1*, *27*), R
41	Nb	1.9 ± 0.2	D–	1.79	1.75	1.58	p	E	D	a?, b, c, h	(*1*, *38*), R
42	Mo	2.16 ± 0.2	D–	2.25	1.99	2.05	p	E	D	b, c, h	(*1*, *27*), R
44	Ru	1.83 ± 0.4	D–	1.93	1.92	1.91	p	E	D	b, c, h	(*1*), R
45	Rh	1.40 ± 0.4	D–	1.25	1.11	1.11	p	E	D	b, c, h	(*1*, *27*)
46	Pd	1.5 ± 0.4	D–	1.76	1.76	1.77	p	E	D	b, c, h	(*27*), R
47	Ag	0.85 ± 0.10	C	1.30	1.30	1.17	p	E	A	c, h	(*86*)
48	Cd	1.85 ± 0.15	B	1.82	1.82	1.73	p	E	A	c, h?	(*34*, *57*, *82*)
49	In	1.65 ± 0.12	C	0.93	0.99	0.65	p, ss	E	A	c, e, g, h	(*57*, *82*, *94*)
50	Sn	2.0 ± 0.4	C–	2.21	2.27	1.88	p	E	B	b, c, h	(*27*, *57*), R
51	Sb	1.0 ± 0.4	D	1.15	1.23	0.91	p	E	D	b, c, h	(*1*, *27*), R
55	Cs	<1.9		1.24	1.22	0.97	ss				(*94*)
56	Ba	2.09 ± 0.11	B	2.23	2.32	2.24	p, ch	E	B–C	a, b, c, e	(*45*, *60*, *75*, *92*)
57	La	1.13 ± 0.3	C	1.30	1.21	1.37	p, ss	E	B	c, e, h	(*26*, *65*, *108*)
58	Ce	1.55 ± 0.2	C	1.58	1.72	1.78	p	E	B	c, h	(*2*, *26*, *78*, *108*)
59	Pr	0.66 ± 0.15	C	0.92	0.88	0.97	p	E	B	c, h	(*26*, *78*, *108*)
60	Nd	1.23 ± 0.3	C	1.54	1.54	1.61	p	E	B	c, h	(*26*, *78*, *108*)
62	Sm	0.72 ± 0.3	C	1.00	1.01	1.11	p	E	B	c, h	(*19*, *26*, *78*, *108*)
63	Eu	0.7 ± 0.3	D	0.58	0.60	0.67	p	E	D	b, c, e, h	(*2*, *37*, *78*)
64	Gd	1.12 ± 0.3	D	1.12	1.20	1.27	p	E	D	b, c, h	(*26*, *78*)
66	Dy	1.06 ± 0.3	D	1.21	1.20	1.30	p	E	D	b, c, h	(*26*, *78*)
68	Er	0.76 ± 0.4	D	1.00	0.99	1.09	p	E	D	b, c, h	(*26*)
69	Tm	0.26 ± 0.2	C	0.18	0.19	0.26	p	E	B	c, h	(*26*, *89*, *108*)
70	Yb	0.9 ± 0.4	D	0.98	0.97	0.99	p	E	D	b, c, h	(*1*, *26*)
71	Lu	0.76 ± 0.3	C–	0.20	0.19	0.13	p	E	B	c, h	(*26*, *108*)
72	Hf	0.8 ± 0.1	C	0.97	1.32	0.99	p	E	B	b, c, h	(*90*, *108*)
74	W	1.7 ± 0.4	D–	0.86	0.85	0.85	p	E	D	b, c, h	(*27*), R
75	Re	≤ −0.3		0.37	0.37	0.42				c, e, h	(*95*)
76	Os	0.7 ± 0.2	D	1.43	1.50	1.54	p	E	D	b, c, h	(*27*, *48*)
77	Ir	0.85 ± 0.2	C	1.51	1.51	1.51	p	E	D	b, c, h	(*16*, *27*)
78	Pt	1.75 ± 0.15	C	1.80	1.76	1.76	p	E	C	b, c, h	(*8*)
79	Au	0.75 ± 0.15	C–	0.96	0.93	0.93	p	E	C	b, c, h	(*85*)
80	Hg	<2.1		1.25	1.32	1.05					(*24*)
81	Tl	0.90 ± 0.17	C	0.93	0.92	0.66	ss	E	B	c, e, g, h	(*56*)
82	Pb	1.93 ± 0.12	B	2.25	2.23	1.86	p	E	B	a,c	(*23*, *40*, *57*, *81*)
83	Bi	<1.9		0.80	0.85	0.93					(*23*)
90	Th	0.2 ± 0.1		(0.30)	0.37	0.32	p	E	B	c, h	(*23*, *40*, *108*)
92	U	<0.60		(−0.36)	0.17	0.10					(*23*)

galaxy, 0.08 to 0.10, valid for the time the sun was formed (*128*).

The iron abundance problem is of quite another character. Here the difficulty is an embarrassment of riches; iron literally dominates the solar spectrum throughout the visible and near-ultraviolet regions. Furthermore, measurements of the *f*-values for the resonance lines of iron are in good accord, but alas we cannot use these strong lines because they are saturated—that is, their profiles are determined largely by damping constants, and good damping constants are not available for iron lines broadened in an atmosphere of atomic hydrogen. On the other hand, for a line that is sufficiently weak that collisional damping is not important (usually a high-level transition), various *f*-value determinations disagree severely with one another.

Earlier photospheric iron abundance determinations (*25, 100, 114*) gave $\log A(\mathrm{Fe}) \sim 6.6$, which not only disagreed with Pottasch's determination (*76*) from the solar corona, but also gave a lower ratio of iron to neighboring elements than would be expected from meteoritic data. The trouble appears to arise from faulty *f*-values for lines that originate in the higher levels. Accordingly, great effort has been expended on the measurement of *f*-values, both to establish absolute values for a limited number of transitions and then to calibrate a large number of relative *f*-values by comparing them with the absolute *f*-values thus established.

Fundamental determinations of *f*-values for subordinate lines are difficult. A number of techniques have been employed: beam-foil (*129*) measurements combined with branching ratios, the electric furnace used as an absorption tube, and the luminous shock tube. Theoretical calculations can, in principle, provide an absolute determination, but comparisons have to be made with experimental results. With weak lines difficulty may be encountered in intermediate coupling calculations.

Numerous methods exist for the measurement of relative *f*-values. One of the most popular is the use of a wall-stabilized arc, burning in argon with a small mixture of some iron compound (*130*). To convert these relative *f*-values to absolute *f*-values, it is necessary to calibrate them with the aid of lines whose absolute *f*-values have already been measured.

A large body of experimental data obtained by different methods (*130, 131*) seems to give a set of mutually consistent *f*-values. These values disagree, however, by an average factor of nearly 3 with those obtained by Gilbert *et al.* (*132*), who measured absolute *f*-values in a glass shock tube system with great care. Although the consensus of current opinion favors the "canonical" *f*-value (*131*) system, and the larger *f*-values measured by Penner's group have *not* been confirmed by other experimentalists, one cannot assert that the last word on the iron *f*-value problem is in. Further fundamental measurements are needed.

The higher iron abundance implied by the lower *f*-values (*131*) is favored by most investigators (*21, 47, 62, 77, 80, 93*). It fits the ratio found for the carbonaceous chondrites and the value found from the solar corona and the solar wind. Independent evidence is found from the forbidden iron lines of the neutral and ionized element, [FeI] and [FeII], which speaks in favor of the higher iron abundance (*29*). The *f*-value question does not enter here. The factor limiting the accuracy here is the weakness of the observed lines and their overlapping or blending with other lines.

From a careful examination of the solar coronal data, Withbroe (*105*) concluded that iron and silicon had nearly the same abundance, which would imply $\log A(\mathrm{Fe}) \sim 7.65$. Smaller values have been proposed by other workers (*17, 63, 71, 98, 99*). The solar flare data, discussed by Crawford *et al.* (*13*), tend to favor the larger value.

To what extent do deviations from local thermodynamic equilibrium play a role? On the basis of their calculations for a 15-level model of FeI, Athay and Lites (*133*) concluded that many FeI lines may be influenced by such departures. Strong lines originating from ground terms and low-lying metastable terms can have strong chromospheric contributions, while high excitation lines may give false abundances because the population of the levels differs from that given by the Boltzmann distribution for a deeper level in the atmosphere. A careful examination of center-limb variations is needed. Except in their cores, some iron lines, however, may be relatively little affected by deviations from local thermodynamic equilibrium.

In conclusion, we feel justified in adopting the iron abundance given in Table 1. The uncertainty ± 0.08 refers to the consistency of individual measurements; the actual error may turn out to be larger than this when the *f*-value problem is finally solved and departures from local thermodynamic equilibrium are fully investigated.

Summarizing Remarks

Except for lithium, beryllium, and boron, the nonvolatile component of the solar atmosphere fits well with data from carbonaceous chondrites. There are a few exceptions, such as indium, but one can be skeptical of the abundance of an otherwise unremarkable metal whose solar abundance is alleged to differ markedly from the meteoritic value. The difference is almost certainly to be attributed to bad *f*-values or to blending or confusion with other lines, or both.

Lithium, beryllium, and boron can be destroyed at the bottom of the solar convection zone. A careful determination of the depletion of these elements will provide a valuable check on models of the past and present sun.

We have not discussed isotope ratios in this article. For all elements that can be investigated from band spectra of their molecules and from isotope effects on hyperfine structure, there is no evidence for any isotope differences between the earth and the sun [for example, see Hauge (*36, 37*)].

There remain, however, a number of pressing problems. We have referred extensively to the need for improved atomic and molecular data, *f*-values, damping constants, collisional parameters needed for non-LTE work, hyperfine structure, and so forth.

On the observational astrophysical side, there is need for improved center-limb measurements of certain critical lines; better data for sunspot spectra; high-resolution spectra at the center, near the limb, and in the ultraviolet beyond the limit 2900 Å imposed by the earth's atmosphere. We need such data for the photospheric spectrum, down to the short wavelengths where it is smothered by the continuous spectrum arising in the chromosphere and corona, but we also need high-resolution spectra of the EUV and x-ray regions. It is hoped that these new, improved data will be obtained and made available as new observations are secured from satellites and rockets.

On the theoretical side, an improved model solar atmosphere incorporating realistic kinematical properties is needed. It will be necessary to develop procedures for handling deviations from local thermodynamic equilibrium in fairly complicated atoms, taking into account the detailed radiation field as it varies with depth in the solar atmosphere. Once reliable *f*-values are secured, the most urgent task will be realistic, accurate handling of non-LTE effects.

For the solar corona it will be necessary to develop models that take into account temporal and spatial variations of temperature, density, and excitation characteristics. The chromosphere will continue to present the greatest challenge of all to both observer and theoretician. To what extent chromospheric line intensities can ever be used in abundance determinations, as for example with helium, is still an open question.

Ultimately, valuable information should be secured from solar wind particles recovered from lunar soils. These data will represent integrations over very long periods of time.

Turning to individual elements, the determination of the solar abundance of helium probably presents the greatest challenge, as it is one of the most important astrophysical composition parameters in construction of models for the sun and in describing the character of the interstellar medium at the time the solar system was formed.

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