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# Abundances of Elements of Cosmological Interest

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# Abundances of elements of cosmological interest

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Observed present-day or proto-solar interstellar mass fractions of relevant species are $X(^1\mathrm{H}) = 0.70$; $X(\mathrm{D}) \approx 3\times10^{-5}$; $X(^3\mathrm{He}) \approx 6\times10^{-5}$; $Y(^4\mathrm{He}) = 0.28$; $X(^7\mathrm{Li}) \approx 5\times10^{-9}$. Because helium is created but not (significantly) destroyed when matter passes through stars, whereas the opposite holds for deuterium, we expect the corresponding primordial values $Y_\mathrm{p}$ and $X_\mathrm{p}(\mathrm{D})$ to be smaller and larger, respectively, than those given above. For deuterium no direct evidence is available yet; theoretically $X_\mathrm{p}(\mathrm{D}) = fX(\mathrm{D})$ where $1.2 \lesssim f \lesssim 5$, depending on details of stellar and galactic evolution. For helium, observations of emission nebulae in irregular galaxies and compact extragalactic H II regions give $Y_\mathrm{p} = 0.23\pm0.01$ (s.e.), consistent with inferences from stellar evolution. Some data on lithium in very old stars suggest that $X_\mathrm{p}(^7\mathrm{Li}) \approx 5\times10^{-10}$.

The most severe constraint on standard three-neutrino Big Bang models comes from $Y_\mathrm{p} < 0.25$, which implies a present-day average baryon density below $2.3\times10^{-31}$ g cm$^{-3}$, compared with a closure density of $1.9\times10^{-29}h_0^2$, where $h_0$ is the Hubble constant in units of 100 km s$^{-1}$ Mpc$^{-1}$. D, $^3$He and $^7$Li are consistent with this limit.

## 1. INTRODUCTION

The nuclear species of cosmological interest are $^1$H, D, $^3$He, $^4$He and $^7$Li (Schramm & Wagoner 1977) and in this paper I shall try to explain what we know about their present abundances in the interstellar medium and what estimates of their primordial abundances we can make when the effects of nuclear synthesis and destruction in stars are allowed for. The interpretation of the resulting constraints on Big-Bang cosmology will largely be left to other papers in this symposium, apart from a small final excursion designed to discuss the mutual consistency of the constraints imposed by the various abundances within the framework of conventional Big-Bang models (Yang *et al.* 1979; Olive *et al.* 1981).

## 2. DEUTERIUM

Ever since the early work of Gamov, deuterium has played an important role in cosmological arguments because its terrestrial and meteoritic abundance is quite large, despite its extreme nuclear fragility. Deuterium is destroyed in stars, rather than created, under most circumstances that can be imagined (see Schramm & Wagoner 1977), whereas the Big Bang provides a natural source if the baryon density is sufficiently low. Because some of the interstellar medium has been 'astrated', that is, has been inside stars and re-ejected with complete destruction of its deuterium, we can assume that the abundance in the interstellar medium either now or at the time of formation of the Solar System is a lower limit to the primordial abundance that can be ascribed to the Big Bang, and this is still one of the most important constraints on cosmology although recently upstaged by $^4$He.

Our knowledge of the abundances of the light elements, including D and $^3$He, has been reviewed by Reeves (1974). Black (1971, 1972) and Geiss & Reeves (1972) estimated the

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proto-solar deuterium abundance on the assumption that a proportion of the $^3$He in the solar wind comes from pre-solar deuterium processed by D(p, $\gamma$) $^3$He. By stepwise heating of gas-rich meteorites, Black discovered a component with a low $^3$He content that he attributed to an ancient solar wind, together with other components having the normal solar wind ratio; subtraction assuming He/H = 0.1 then leads to

$$\mathrm{D/H} = (2.5 \pm 1.0) \times 10^{-5} \quad \text{or} \quad X_2 = 3.5 \times 10^{-5}.$$

A similar value is deduced from an absorption line of HD in Jupiter, and it follows that deuterium in terrestrial and meteoritic water is overabundant by a factor of about 6 owing to fractionation.

In 1972, deuterium was also detected for the first time in the interstellar medium by radio observations, particularly of DCN and other molecules, and HD absorption lines were found in the ultraviolet with the Copernicus satellite (Spitzer & Jenkins 1975). Owing to fractionation, again, these observations are not very useful for quantitative abundance determinations, but Rogerson & York (1973) discovered atomic lines of deuterium on the line of sight to β Centauri and since then D/H ratios have been measured from interstellar lines observed with Copernicus in about a dozen stars at distances up to nearly 1 kpc (Laurent *et al.* 1979 and references therein; York 1982) with varying degrees of precision, depending on the complexity or otherwise of the velocity field resulting from multiplicity of clouds in the line of sight. In the best cases the uncertainty is about 30%, but despite this the measured values range over more than an order of magnitude. Bruston *et al.* (1981) attribute most of these variations to the effects of radiation pressure and molecular fractionation, and they estimate the most probable value of the abundance to be

$$\mathrm{D/H} = (2.5 \pm 0.25) \times 10^{-5} \quad \text{or} \quad X_2 = 3.5 \times 10^{-5},$$

which is very close to the estimates for the proto-solar system and thus probably reliable within a factor of 2 or better.

To estimate the pregalactic value, it is necessary to correct for astration, and the result of this procedure is highly model-dependent. Audouze & Tinsley (1974) have given numerically computed evolutionary abundance curves for deuterium and other light elements on the basis of specific galactic evolutionary models, but it is of interest to develop a simple analytical theory based on the instantaneous recycling approximation because this gives a more direct insight into the effects of different parameters assumed in the models. (The instantaneous recycling approximation assumes that all the astration effects occur in negligible time compared with the timescale of galactic evolution.)

We assume that evolution takes place in a system having a mass $m(t)$ consisting of interstellar gas with mass $g(t)$ and stars with total mass $s(t)$ increasing with time, i.e.

$$m(t) = g(t) + s(t), \tag{2.1}$$

with no stars at the beginning, i.e.

$$s(0) = 0, \tag{2.2}$$

and with the deuterium abundance $X_2$ in the gas having initially its primordial value

$$X_2(0) = X_{\mathrm{p}2}. \tag{2.3}$$

We assume, further, that, whenever a generation of stars is formed, a constant fraction $\alpha$ remains locked up in long-lived stars of low mass or compact remnants such as white dwarfs and neutron stars and the remainder, $(1-\alpha)$, is returned to the interstellar medium with its deuterium destroyed. Unprocessed material with the primordial deuterium abundance $X_{\mathrm{p}2}$ arrives from the intergalactic medium at some rate $\mathrm{d}m/\mathrm{d}t$. The abundance of deuterium in the interstellar gas is then governed by the differential equation

$$\frac{\mathrm{d}}{\mathrm{d}s}\,(X_2 g) = -\frac{X_2}{\alpha}+X_{\mathrm{p}2}\,\frac{\mathrm{d}m}{\mathrm{d}s}, \tag{2.4}$$

with the initial conditions (2.2) and (2.3). It is of interest to look at the solutions of (2.4) for three different sorts of galactic evolutionary model (cf. Tinsley 1980; Pagel 1981).

(i) A simple closed model with $\mathrm{d}m/\mathrm{d}s = 0$. Then

$$X_2/X_{\mathrm{p}2} = (g/m)^{(1-\alpha)/\alpha}. \tag{2.5}$$

Such a model is equivalent for this purpose to one that I have advocated for the solar neighbourhood (Pagel 1981), based on models of dynamical collapse by Larson (1976) and Lynden-Bell (1975). The ratio $g/m$ is probably between 0.1 and 0.25 (Tinsley 1981) and $\alpha$ between 0.6 and 0.8 (Pagel 1981), whence

$$0.2 < X_2/X_{\mathrm{p}2} < 0.7,$$

which indicates the wide range of uncertainty that exists.

(ii) An extreme inflow model (Larson 1972) in which star formation is balanced by inflow and $g(t) = \text{constant}$, $\mathrm{d}m/\mathrm{d}s = 1$. In this case,

$$X_2/X_{\mathrm{p}2} = \alpha+(1-\alpha)\,\mathrm{e}^{-[(m/g)-1]/\alpha} \rightarrow \alpha, \tag{2.6}$$

and the depletion is only by a factor of $0.7\pm 0.1$. This model, however, is only of academic interest because it does not predict the observed increase in stellar metallicity with time (Pagel & Patchett 1975; Twarog 1980).

(iii) An intermediate model favoured by the Yale astronomers (Twarog 1980; Tinsley 1981) in which inflow occurs at a smaller rate such that

$$\mathrm{d}m/\mathrm{d}s = \lambda,$$

where $\lambda$ is a constant of the order of 0.3. For this model,

$$\frac{X_2}{X_{\mathrm{p}2}} = \frac{\alpha\lambda}{1-\alpha'}+\frac{1-\alpha}{1-\alpha'}\left(\frac{g}{\lambda g+(1-\lambda)\,m}\right)^{(1-\alpha')/\alpha'} \rightarrow \frac{\alpha\lambda}{1-\alpha(1-\lambda)}, \tag{2.7}$$

where $$\alpha' \equiv \alpha(1-\lambda),$$

giving finally $$0.37 < X_2/X_{\mathrm{p}2} < 0.73,$$

so that a factor of 2 depletion is a reasonable estimate, but the uncertainties are again substantial. The conclusion, then, is that

$$X_{\mathrm{p}2} \approx 5\times 10^{-5}$$

with an uncertainty by a factor of 3 or so.

There is one approach that may, in time, lead to a more accurate estimate of the primordial deterium abundance. Sargent *et al.* (1980) have identified the 'Lyman only' absorption-line systems observed in quasars of high red shift (Lynds 1971) with primordial hydrogen clouds consisting of unprocessed and therefore unastrated material. In suitable objects, it should be possible to make a search for deuterium lines and thereby obtain interesting limits for primordial deuterium direct.

## 3. HELIUM-3

Knowledge of the abundance of $^3$He is quite limited. Black's work and related arguments lead to a proto-solar abundance (see Reeves 1974)

$$^3\mathrm{He/H} = (2 \pm 1) \times 10^{-5} \quad \text{or} \quad X_3 = 4 \times 10^{-5}.$$

The 3.46 cm hyperfine line of $^3\mathrm{He}^+$ has been tentatively detected in the H II region W51 with an intensity corresponding to twice this abundance (Rood *et al.* 1979). Ulrich (1971) suggested a 10 times higher value on the basis of a tentative identification of a $^3$He-burning sequence in young star clusters, but the balance of the evidence evidently favours something in the neighbourhood of Black's upper limit, and the problem is how to correct it for galactic chemical evolution.

This question has been discussed by Rood *et al.* (1976) who calculate that astrated material that has been through low-mass stars and re-ejected from them in planetary nebulae is enriched to $X_3 \approx 10^{-3}$ as a result of dredge-up from hydrogen-burning zones. Consequently, if as little as 30 % of the interstellar gas has been astrated, which is near our lower limit calculated for deuterium, then this gas as a whole should have $X_3 \geqslant 3 \times 10^{-4}$, over and above any cosmological value, and our problem is to explain why so *little* $^3$He is observed. Rood *et al.* conclude from this that, even if the problem could be solved by bridging the various uncertain parameters, any cosmological contribution to $^3$He would be entirely swamped by stellar nucleosynthesis (cf. Tinsley 1977). However, the likelihood that astration will increase $^3$He rather than destroy it permits some interesting cosmological conclusions to be drawn by taking the proto-solar value as an upper limit (Schramm, this symposium; Yang *et al.* 1982). If, on the other hand, we understand stellar evolution so badly that stars actually destroy $^3$He without our realizing the fact, then only a much less stringent upper limit can be placed on $X_\mathrm{p}(^3\mathrm{He})$ by the same sort of considerations as were used for deuterium, giving

$$X_\mathrm{p}(^3\mathrm{He}) \leqslant 3 \times 10^{-4}.$$

This question is discussed further by Schramm (this symposium).

## 4. HELIUM

### 4.1. *General comments on the abundance of helium*

Helium is not only the second most abundant element in the visible Universe, but also the one about which we have the most knowledge. Its abundance can be derived from absorption lines in hot stars, from emission lines in gaseous nebulae and from global stellar properties: the mass–luminosity relation, pulsation characteristics and location of the zero-age or evolved main sequence in the luminosity-effective temperature plane; and also from certain statistical properties like the relative numbers of red giants and horizontal-branch stars in globular clusters. These methods do not all have the same precision, but in favourable cases (notably H II regions) the precision can be better than 10 % and possibly even 5 %. This contrasts with the situation for other elements in astrophysical objects, where a precision of 20 % can be rarely, if ever, surpassed. Generally, the abundance of helium, by mass, in objects not greatly affected by internal nuclear processing, is found to be in the range $0.2 < Y < 0.3$, which is within the range that is of interest for cosmology (Schramm & Wagoner 1977). The questions that arise, then, are:

(i) Is there *any* object that may lead us to doubt the existence of a universal pre-galactic helium abundance of the order of 0.2 or more as required by canonical Big-Bang models? and

(ii) Can we discern a trend of helium abundance with the mass-fraction $Z$ of carbon and heavier elements synthesized in stars that will enable us to sort out the contribution of stellar nucleosynthesis to the helium in any particular class of objects and thereby extract a primordial value with sufficient precision to provide interesting constraints on current cosmological models?

The answer to the first question has been generally considered for at least 10 years to be definitely no, for reasons that were succinctly summarized by Searle & Sargent (1972*a*). Data available at that time (see, for example, Danziger 1970) did indicate a few apparent possible exceptions to a universal helium abundance of at least 0.2 by mass, notably (i) a supposed weakness of helium lines in quasars (but see Williams 1971), (ii) weak helium lines among the so-called subdwarf B stars in globular clusters and among field stars in the galactic halo, and (iii) the large mass deduced from a supposed detection and separation measurement of a faint companion to the nearby astrometric binary subdwarf μ Cassiopeiae (Hegyi & Curott 1970; but see Faulkner 1971). Actually this last result now appears to have been spurious because when the companion was eventually found later on (by a photographic technique) it turned out to be much fainter than Hegyi & Curott could have detected and there is no evidence for low helium in μ Cas (Feibelman 1976). The same applies to the similar subdwarf system 85 Pegasi (Smak 1960; Feierman 1971) although in both cases the uncertainties are still considerable. The results for the subdwarf B stars, by contrast, are not wrong but they are misleading in so far as they refer only to atmospheric layers, which are subject to gravitational settling (Greenstein *et al.* 1967). Sargent & Searle (1967) provided the first proof that the atmospheric layers are unrepresentative by discovering phosphorus lines in the halo B subdwarf Feige 86, which thus shows similarities to the young chemically peculiar star 3 Cen A, which it also resembles in having an abnormally high ratio of $^3$He/$^4$He (Hartoog 1979); similar conclusions have since then been extended to several other B subdwarfs when observed with sufficiently high resolution (Greenstein & Sargent 1974). However, by the same token we should remain unimpressed by the *normal* helium abundances found in various halo objects in an advanced evolutionary state where they can have been affected by hydrogen burning, e.g. the luminous B star Barnard 29 in the globular cluster M13 (Auer & Norris 1974) and the planetary nebula K648 in M15 (Hawley & Miller 1978). A further signal contribution made by Searle & Sargent to this issue was the discovery that two of Zwicky's compact galaxies, I Zw18 and II Zw40, have spectra characteristic of H II regions with substantial deficiencies of heavy elements (O and Ne) but near-normal helium (Sargent & Searle 1970; Searle & Sargent 1972*b*). Meanwhile, further studies of quasars have shown that, while He$^+$ λ4686 Å is often weak, the He I lines are in fact remarkably strong, which is explained by fairly intricate radiative transfer processes combined with a more or less normal helium abundance (Davidson & Netzer 1979). Consequently there is no evidence that old objects exist with $Y < 0.2$ or so.

### 4.2. *Helium in the solar neighbourhood*

Before examining whether we can find a correlation between helium and heavy-element abundances that can be extrapolated to pre-galactic or pre-stellar helium, we should first satisfy ourselves that there is indeed a unique helium abundance in some astrophysical system at a given time, and this is something that can be tested by looking at the Sun and at stars

Table 1. Helium abundance in the Sun and its neighbourhood

| object | method | $y$ | $Y$ | references |
|---|---|---|---|---|
| | *Sun* | | | |
| solar interior, envelope ($Z = 0.02$) | evolution models; oscillation frequency | $0.085 \pm 0.009$ | $0.25 \pm 0.02$ | 1 |
| prominences | emission lines | $0.100 \pm 0.025$ | $0.28 \pm 0.05$ | 2 |
| flare particles | plastic emulsion, etc. | 0.06: | 0.19: | 3 |
| solar wind | counters | $0.04 \pm 0.01$ | $0.14 \pm 0.03$ (fractionation) | 4 |
| | *young objects* | | | |
| B, A main sequence interiors | binary mass–luminosity relation | $0.11 \pm 0.02$ | $0.28 \pm 0.04$ | 5 |
| normal B stars: atmospheres | absorption-line profiles in blue | $0.10 \pm 0.02$ | $0.28 \pm 0.04$ | 6 |
| Orion nebula | optical recomb. lines | $0.100 \pm 0.005$ | $0.28 \pm 0.01$ | 7 |
| | radio recomb. lines | $\geqslant 0.088$ | $\geqslant 0.26$ | 8 |

*References:* 1, Bahcall & Ulrich (1971), Gough & Weiss (1976), Christensen-Dalsgaard & Gough (1980); 2, Heasley & Milkey (1978); 3, Crawford *et al.* (1975), assuming $12 + \lg(\mathrm{O/H}) = 8.9$; 4, Hundhausen (1972); 5, Popper *et al.* (1970); 6, Leckrone (1971); Norris (1971), Auer & Mihalas (1971); 7, Peimbert & Torres-Peimbert (1977); 8, Thum *et al.* (1980), Thum (1981).

and H II regions within a few hundred parsecs. Planetary nebulae are to be excluded because in some cases at least they are appreciably self-enriched in helium (Torres-Peimbert & Peimbert 1977), and we should also note that the solar value could in principle be slightly different from that found in younger objects because of time-dependence and large-scale spatial gradients in the enrichment. A variety of data, based on several different techniques, are available for the solar neighbourhood (table 1) and the results in the table are mutually consistent if we assume that solar flare particles and the solar wind are affected by selective acceleration processes. However, since the determination from the solar interior could be suspect until the solar neutrino problem has been solved, the initial abundance of helium in the Sun largely rests on the theory of radiative transfer in prominences, and the result is still only of low precision. The results for hot stars that show absorption lines of helium in their optical spectra are also of rather low precision because of small uncertainties arising from line-broadening theory and departures from local thermodynamic equilibrium. Probably the most accurate value comes from the study of recombination lines of hydrogen and helium in the Orion nebula, in which three of the helium lines ($\lambda\lambda$ 4471, 5876, 6678 Å) have been shown to depart from optically thin, purely radiative conditions by only about 5% (Peimbert & Torres-Peimbert 1977). A somewhat larger source of uncertainty is the correction for ionization, since the recombination lines actually give the ratio $y^+$ of singly ionized helium to ionized hydrogen, weighted by electron density in the line of sight. This leads to a position-dependent ionization correction factor (i.c.f.) to allow for neutral helium, which can be estimated in optical observations by looking at different areas and comparing with the $\mathrm{O}^+/\mathrm{O}^{2+}$ ionization balance (the i.c.f. is virtually 1.00 close to the Trapezium). For radio data, the i.c.f. is more difficult to determine because of the large beam width and because other ions are not available, and for this reason only a lower limit is quoted. Doubly ionized helium, when present, is readily detected and allowed for in the course of optical observations; it is absent or very weak in normal H II regions.

Many nearby stars show unusually high or low helium abundances in their photospheres. Large helium abundances can result from hydrogen burning, mixing and mass loss in the course of stellar evolution: examples are Wolf–Rayet stars, hydrogen-deficient carbon stars and their hotter counterparts known as helium stars (Hunger 1975; Baschek 1979), the DB white dwarfs (Weidemann 1975) and some planetary nebulae (Torres-Peimbert & Peimbert 1977; Kaler 1978). Such effects are, of course, entirely in line with what one would expect from stellar nucleosynthesis. More 'anomalous' are stars with unusually low atmospheric helium, which include not only the DA white dwarfs (Weidemann 1975) but also the large group of so-called chemically peculiar (Bp, Ap and Am) stars, which are main-sequence stars with effective temperatures between 7000 and 20000 K or so (Preston 1974). The photospheres of these stars, some of which have strong magnetic fields up to 1 T, show an extraordinary variety of abundance anomalies in which rare elements such as phosphorus, manganese, rare earths and mercury are enhanced, as are somewhat less rare elements like silicon and chromium, while abundant elements like helium and oxygen are deficient, with different sorts of anomalies occurring predominantly in different ranges of effective temperature. Most are relatively slow rotators and many are spectrum variables in which the abundance anomalies vary as different faces of the star are presented to the observer, which is one of several reasons for believing that the abundance anomalies are only a skin effect, not representative of the composition of the star as a whole; the most successful theory to account for them is that of diffusion, i.e. gravitational settling modified by selective radiation pressure (Michaud 1970; Vauclair & Vauclair 1979). Thus in relatively stable atmospheres with no deep convection zone, only weak turbulence caused by rotation, and possibly a magnetic field to inhibit mixing currents, helium sinks down below the atmosphere when the effective temperature is below about 20000 K, while at higher temperatures it can be pushed up by radiation pressure in the ionization continuum and stars then may appear superficially helium-rich, although in some cases they are probably truly helium-rich as a result of nuclear processing (Hunger 1975). Naturally all such chemically peculiar individual stars should be disregarded in assessing the helium abundance in the solar neighbourhood, which then comes out at $Y = 0.28 \pm 0.02$. However, significantly lower values have been reported from photoelectric spectrophotometry of ten or more main-sequence stars in h and $\chi$ Persei and the association Cepheus OB III (Nissen 1976) and in the nearby Hyades cluster (Strömgren *et al.* 1982); if these results cannot be somehow 'explained away' then we shall have to accept the possibility of quite large variations in helium abundance at the present time.

### 4.3. *Helium at other places and other times*

We assume that massive stars synthesize both helium and heavy elements and supply them to the interstellar medium. A simple, naïve model of galactic chemical evolution akin to model (i) in §2 then leads one to expect a linear relation

$$Y = Y_p + Z\mathrm{d}Y/\mathrm{d}Z \tag{4.1}$$

between helium and heavy element abundances to hold at all places and times, if one ignores possible differences in past rates of formation of stars of different mass, exchanges of material with the surroundings and other complicating factors. The range of strict applicability of this equation is quite narrow, but it has the saving grace that, when $Z$ is sufficiently small, we can still obtain interesting results from the more general assumption that $Y \to Y_p$ from above.

To extrapolate to $Z = 0$, we can go back in time by studying nearby metal-deficient objects (subdwarfs) and globular clusters, which latter may also involve spatial gradients across the

Galaxy. Alternatively we can study luminous young objects widely distributed in space, e.g. H II regions in our own and other galaxies in which the heavy-element abundances are different, basically because of the relative youth of the underlying stellar populations.

### (i) *Globular clusters and field stars of the galactic halo population*

In old stellar populations, the stars still on the main sequence are too cool to show helium lines in their spectra, but the theory of stellar structure predicts a relation between the position of the main sequence in the colour–luminosity (Hertzsprung–Russell (H–R)) diagram and the abundance parameters $Y$ and $Z$ in the sense that increasing $Z$ pushes the sequence up and to the right while increasing $Y$ does the opposite. In this way Carney (1979) has used the difference of the order of half a magnitude in luminosity between extreme subdwarfs and main-sequence stars of the Hyades cluster to estimate an average value of $Y$ in the subdwarfs. This result (table 2) is consistent with those derived earlier on similar grounds by Cayrel (1968), Dennis (1968) and Eggen (1973). However, similar differences are expected between nearby field stars differing in metallicity from the Sun by quite small factors of 2 or 3, and these are not clearly perceived (Eggen 1973; Perrin *et al.* 1977). A strict cancellation of main-sequence shifts can be estimated to result for such stars if $\mathrm{d}Y/\mathrm{d}Z \approx 5$, but most, if not all, of this factor can be accounted for by inadequacies in the data, especially parallaxes (Cayrel 1978).

Table 2. Helium abundances in old stars

| | globular clusters | | | | field stars | |
|---|---|---|---|---|---|---|
| | M 92 | M 3 | M 13 | 47 Tuc | subdwarfs | references |
| [Fe/H]† | −2.3 | −1.8 | −1.5 | −0.9: | −1.7 | |
| $Z$ | 0.0002‡ | 0.0005‡ | 0.0005 | 0.002: | 0.0008‡ | |
| $T$ (RR blue edge) | $0.22 \pm 0.07$ | $0.22 \pm 0.07$ | | | | 1 |
| $\Delta T$ (instab. strip) | $0.29 \pm 0.04$ | $0.29 \pm 0.04$ | | | | 2 |
| RR periods | 0.3 | 0.3 | | | | 3 |
| HB/RG ratio | | $0.23 \pm 0.04$ | | $0.24 \pm 0.04$ | | 4 |
| H–R diagram | $0.26 \pm 0.03$ | $0.30 \pm 0.03$ | $0.30 \pm 0.03$ | $0.26 \pm 0.03$ | | 5 |
| main sequence | | | | | $0.23 \pm 0.04$§ | 6 |

† $[X] \equiv \lg X\,(\text{object}) - \lg X\,(\text{Sun})$.
‡ Assuming [O/Fe] = 0.5.
§ Assuming $Y = 0.30$ in the Hyades.
*References:* 1, Sandage (1969), revised after Tuggle & Iben (1972); 2, Deupree (1977); 3, Castellani (1981); 4, Renzini (1977), cf. Buonanno *et al.* (1981); 5, Demarque & McClure (1980), Caputo & Cayrel de Strobel (1981) (adapted); 6, Carney (1979).

Several methods exist for the determination of helium abundance in globular clusters, but these are subject to rather uncertain systematic errors due to theoretical modelling of stellar structure and evolution and to uncertainties in the abundances of C, N and O. Metal abundances are still uncertain by factors of about 2, oxygen is usually deficient by smaller factors than the metals, and C and N vary, often within a single cluster, by large amounts that may be partly primordial in origin (see Kraft 1979; Pagel & Edmunds 1981; Freeman & Norris 1981). The methods used involve one of the following: (i) effective temperatures and periods of pulsating variables (RR Lyrae stars); (ii) morphology of the H–R (luminosity–colour) diagram; and (iii) the population ratio of horizontal-branch stars to red giants. Helium abundances differing by small amounts of the order of 0.03 from cluster to cluster also constitute one of the many candidates for the 'second parameter' that is postulated in order to account for the existence of differing horizontal-branch morphologies in different clusters with the same metallicity (Kraft 1979; Freeman & Norris 1981), but not necessarily the most promising one. A number of estimates of helium abundance in a few well studied globular clusters are collected in table 2.

Pulsation theory (Christy 1966) provides a useful method of estimating helium abundances in stellar envelopes in those clusters where there are sufficient variables to define the boundaries of the instability strip, the basic effect being the importance of $He^+$ as a driver of pulsation. The results are relatively insensitive to mass, age and heavy-element abundance, within reasonable limits, but they do depend critically on opacities and on the exact determination of effective temperature or of its dependence on colour. Alternatively, if one can predict the width of the instability strip (Deupree 1977) or the period distribution of RR Lyrae stars (Castellani 1981), then the observational uncertainties are reduced at the expense of placing heavier reliance on convection theory or evolutionary stellar models.

The effect of varying initial helium abundance on the morphology of the H–R diagram is that, at a given heavy-element abundance $Z$, increasing $Y$ pushes the unevolved main sequence down and the 'zero-age' horizontal branch up. As a result, the relative luminosities of horizontal-branch and main-sequence stars are altered by nearly 0.4 dex when $Y$ changes from 0.2 to 0.3. Such differences are indeed found among the H–R diagrams, leading to a suggestion that there could be significant differences between different globular clusters, and that they may not be in a one–one correspondence with their metal contents in all cases (Demarque & McClure 1980; Caputo & Cayrel 1981; see also Sandage *et al.* 1981). This last result, however, is quite critically dependent on the heavy-element contents, particularly for 47 Tuc, where further revisions might well be enough to bridge the gap.

The third method of estimating helium abundances in globular clusters relies on the fact that, when there is more helium, the luminosity and the lifetime on the horizontal branch are both increased, while the lifetime on the first ascent of the red-giant branch above any given luminosity is decreased. Consequently the ratio of the number of horizontal-branch stars to that of red giants more luminous than the horizontal branch is an indicator of helium abundance (Iben 1968), although one that is subject to various theoretical and experimental problems, chiefly semi-convection and completeness respectively (Renzini 1977). Finally, Green (1981) has made a careful study of the luminosity functions of the red giants in many globular clusters and compared them with theoretical luminosity functions for $Y = 0.2$ and 0.3. In most cases the agreement is good but, for the four clusters less than 6 kpc from the galactic centre that were studied, she found a significant deficiency of stars near the tip of the red giant branch compared with what was expected on the basis of a constant horizontal-branch luminosity for all clusters. Relaxing this assumption leads to a tentative suggestion that the horizontal branch is brighter than expected for these inner clusters, either because of an enhanced core mass (e.g. due to rotation) or because of enhanced helium abundance, i.e. $Y > 0.3$.

The end result of the data summarized in table 2 (and similar data for several other globular clusters) is that the uncertainties just about cover the range of interest, which is encouraging in a way, but is of no help in estimating $Y_p$.

### (ii) *Galactic and extragalactic* H II *regions*

Our remaining hope of setting interesting limits on $Y_p$ (and $dY/dZ$) is provided by observations of hydrogen and helium recombination lines in H II regions, i.e. clouds of interstellar material ionized by ultraviolet radiation from one or more hot stars newly formed from the same cloud or cloud complex. These can be H II regions in our own Galaxy, where the heavy element fraction $Z$ decreases outwards from the centre at a rate of 0.05 dex per kiloparsec or so (see Pagel & Edmunds 1981); or in the Magellanic Clouds where $Z$ is significantly lower than in the solar neighbourhood; or giant H II regions in external spirals that have similar or in some cases larger abundance gradients than are found in our own Galaxy; or gas-rich external galaxies discovered in surveys such as those by Zwicky and Markarian, which have emission-line spectra typical of H II regions and cover a wide range of heavy-element abundance from about 0.02 of the solar value upwards and which are referred to variously as extragalactic H II regions and as non-Seyfert emission-line galaxies. Many studies of abundances in all these classes of H II regions have been made (see Pagel & Edmunds 1981), but the determination of $Y_p$ makes such exceptional demands on precision of observations and data analysis that only a handful of investigations, undertaken specifically with this purpose in mind, need to be considered here.

An early attempt to correlate $Y$ and $Z$ and extrapolate to $Z = 0$ was made by Peimbert & Torres-Peimbert (1974, 1976) in studies of the two Magellanic Clouds, which have the advantage of having many H II regions, which are easily resolved so that the i.c.f. can be studied point by point, and a more or less uniform abundance in each cloud (Pagel *et al.* 1978). By joining the two clouds to Orion and assuming equation (4.1), they deduced $Y_p = 0.228 \pm 0.014$ and $dY/dZ \approx 2.7$, but the result is open to question because of having only three data points and because the chemical evolution of the solar neighbourhood including Orion may be quite distinct from that of Irregular galaxies so that (4.1) need not apply to it. This reservation was overcome in a subsequent investigation by the Peimberts and colleagues (Lequeux *et al.* 1979) in which the Magellanic Cloud data were combined with new data for six Irregular and compact galaxies from which they deduced

$$Y = 0.233 \pm 0.005 + (1.7 \pm 0.9)\ Z, \tag{4.2}$$

where $Z$ is taken to be proportional to the oxygen abundance. (Inclusion of Orion increases $dY/dZ$ to 2.8 while leaving $Y_p$ virtually unaffected at 0.228.) The large $dY/dZ$ ratio was explained by computing galactic enrichment models in which massive stars are assumed to undergo heavy mass loss in the core hydrogen and helium burning stages (Chiosi & Caimmi 1979; Chiosi & Matteucci 1982), which basically reduces the heavy-element yield while leaving the helium yield unaltered (see Serrano & Peimbert 1981). However, Maeder (1981) has challenged the theoretical basis of these computations.

Since then, there have been some other investigations reaching conclusions that agree with those of Lequeux *et al.* (1979). Thum (1981) has discussed the optical and radio recombination line data in our own Galaxy and suggests that there is a helium abundance gradient corresponding to the $Z$ gradient, although its precise run is rendered somewhat uncertain by the problem of ionization corrections. Assuming that $dY/dZ \approx 2$, Thum deduces

$$Y_p = 0.22 \pm 0.02,$$

but it is fair to say that the whole question of helium in galactic H II regions remains to be settled by further studies (Shaver *et al.* 1982). Recently Rayo *et al.* (1982) have carried out an extremely careful study of H II regions in the Scd spiral M101 in which they allowed for underlying stellar absorption and emission lines as well as radiative transfer effects and obtained helium abundances for two H II regions in which they were satisfied that they had adequate ionization corrections and again found a helium gradient decreasing outwards. From the outer H II region, NGC 5471, they derive the pre-galactic value

$$Y_p = 0.216 \pm 0.010(3\sigma) + 0.004(1.7 - \mathrm{d}Y/\mathrm{d}Z),$$

which is certainly the highest precision ever claimed for an estimate of $Y_p$.

Other investigations of this problem have led to considerably less clearcut results, and in some cases cast doubt on the whole idea of a straightforward single-valued relation between $Y$ and $Z$. Samples of Irregular galaxies and extragalactic H II regions studied by French (1980), Talent (1980), Kinman & Davidson (1981) and Kunth (1982) yield essentially a scatter diagram in which no correlation between $Y$ and $Z$ can be discerned unless one includes Orion, which is not permissible. If indeed the relation is a scatter diagram, then the most logical way to proceed is simply to take a straight mean of the helium abundances as an upper limit, which is what Kunth does, leading to

$$Y_p \leqslant 0.245 \pm 0.003\ (1\sigma),$$

which is still an interesting result.

Why is there so much disagreement between the results of Lequeux, Peimbert and their colleagues and those of the other workers? The problem is that the $\mathrm{d}Y/\mathrm{d}Z$ effect is so small as to stretch the precision of the data to the limit, and it may be that the other investigations simply do not have enough precision. Alternatively, there could be a systematic effect in the ionization correction factors of Lequeux *et al.* Finally, there may be a real scatter and Lequeux *et al.* could just have been lucky (or unlucky) in their choice of objects. This point is investigated in table 3, where we examine the precision of the data both as claimed by the authors and as emerges from the standard deviation relative to the simple assumption that $y$ is constant. It is clear that the errors in all except Lequeux's and Kunth's data are too large to reveal any correlation even if it exists. The two latter data sets have comparable errors, with Kunth's data showing a slightly larger scatter; there is only one object in common, II Zw 40, for which the agreement is excellent. These two data sets are shown in figure 1, together with weighted least-squares regressions and their $1\sigma$ errors; the two relations are not mutually inconsistent, but the uncertainties in the $\mathrm{d}Y/\mathrm{d}Z$ slope are evidently large. I conclude that the best estimate that can be made of the primordial helium abundance at present is

$$Y_p = 0.23 \pm 0.01\ (\text{s.e.}),$$

and that we can set a firm upper limit of

$$Y_p < 0.25,$$

which is no different from the conclusions of Yang *et al.* (1979). To establish that the true value is as low as 0.22 would, of course, provide a still more interesting constraint on cosmological models, but no firm conclusion can be reached on this point until we have more numerous data of very high precision.

TABLE 3. $Y$ AND $Z$ IN IRREGULAR GALAXIES AND EXTRAGALACTIC H II REGIONS

| | Lequeux *et al.* | Kunth | French | Talent | Kinman & Davidson |
|---|---|---|---|---|---|
| range in $Z$† min | 0.0004‡ | 0.0008 | 0.0004 | 0.0018 | 0.0004 |
| max | 0.0063‡ | 0.0049 | 0.0050 | 0.0064 | 0.0069 |
| number of objects | 10 | 13 | 10§ | 26 | 9 |
| mean $y$ | 0.080 | 0.082 | 0.087 | 0.077 | 0.096 |
| r.m.s. error (authors') | 0.005 | 0.0053 | 0.010 | 0.016 | 0.028 |
| sample standard deviation | 0.0038 | 0.0054 | 0.016 | 0.009 | 0.031 |
| $\bar{Y}$ | 0.242 | 0.246 | 0.257 | 0.235 | 0.277 |
| s.e.m. | ± 0.003 | ± 0.003 | ± 0.012 | ± 0.004 | ± 0.024 |

† Assuming $Z = 25$ (O/H).
‡ Temperature fluctuations taken to be zero in computing $Z$.
§ Low-luminosity galaxies only.

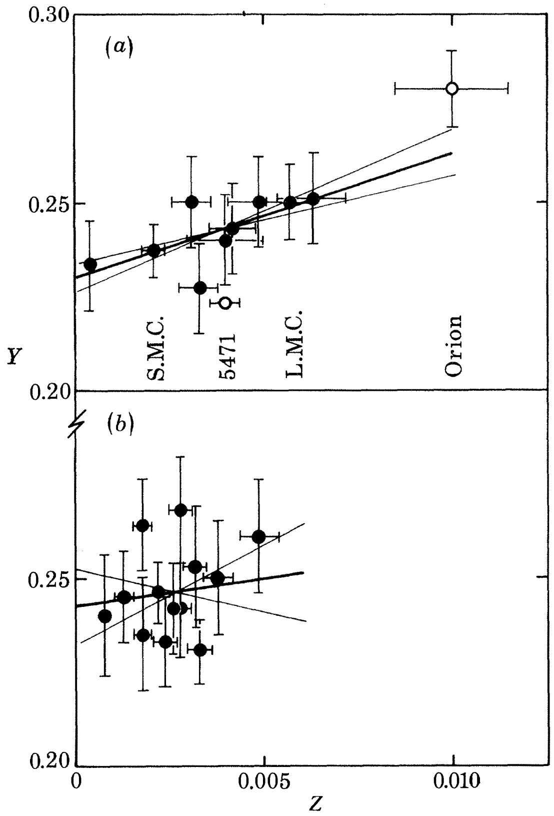

FIGURE 1. Relation between $Y$ and $Z$ in extragalatic H II regions, based on the results of (*a*) Lequeux *et al.* (1979) and (*b*) Kunth (1982). It has been assumed that $Z = 25(\mathrm{O/H})$ and that temperature fluctuations in the H II regions are negligible. The least-squares regressions have the equations $Y = 0.230 \pm 0.004 + (3.31 \pm 0.95)\ Z$ (Lequeux *et al.*) and $Y = 0.243 \pm 0.010 + (1.4 \pm 3.8)\ Z$ (Kunth). Orion and NGC 5471 are shown with the data of Lequeux *et al.* for comparison; they are not included in the solution. Error bars are $\pm 1\sigma$.

## 5. LITHIUM-7

The abundance of $^7$Li in meteorites corresponds to $X_7 = 5\times 10^{-9}$ (Cameron 1981), but in stellar atmospheres lithium is usually depleted by an amount that increases with the age of the star and with the depth of the surface convection zone because it is destroyed by hydrogen-burning at the relatively low temperature of $2\times 10^6$ K (see review by Boesgaard 1976). The abundance in the youngest stars is consistent with that in meteorites, while in the interstellar gas lithium is depleted by a factor similar to sodium and potassium, probably through being locked on grains (see Snell & Vanden Bout 1981). On the other hand, very large lithium abundances, up to 100 times the cosmic value, are found along with excesses of carbon and s-process elements in certain cool carbon stars like WZ Cas and WX Cyg, which is accounted for by the synthesis of its radioactive progenitor $^7$Be through $^3$He$(\alpha, \gamma)$ $^7$Be (Cameron & Fowler 1971) in the interior, followed by dredge-up to the atmosphere when the outer convection zone penetrates between helium shell flashes (Sackmann *et al.* 1974). This process, together with others such as low-energy cosmic rays, probably leads to a net enrichment of $^7$Li in the present-day interstellar medium over the pre-galactic value (Audouze & Tinsley 1974), which can therefore only be estimated if one can find very old stars that at the same time have sufficiently thin convective envelopes for surface lithium not to have been substantially depleted during their lifetime.

Recently Spite & Spite (1981, 1982) have discovered lithium lines in the spectra of five extreme subdwarfs deficient in metals by factors between 15 and 250 and with effective temperatures between 5000 and 6000 K, with the coudé spectrograph of the new Canada–France–Hawaii Telescope. The interesting features of their results are that (i) lithium is actually quite abundant, with about one tenth of the cosmic value, and (ii) the abundance found is the same in the four hotter stars, independent of their temperature or metallicity, and significantly lower only in the coolest star (Groombridge 1830) of the sample, which they take as tentative evidence that lithium destruction in the outer convection zone may not have been effective in these four stars. (This contrasts with the situation in old metal-rich dwarfs, where there is a systematic gradual decrease in surface lithium abundance with decreasing effective temperature.) Spite & Spite conclude that they may well have identified the pre-galactic lithium abundance with

$$X_p(^7\text{Li}) \approx 5\times 10^{-10},$$

and an error of perhaps a factor of 3. In any case they have established an interesting lower limit.

## 6. SUMMARY AND CONCLUSIONS

Our estimates of present-day and primordial abundances of the relevant elements are summarized in table 4 and plotted against the predicted Big-Bang values taken from Yang *et al.* (1979), assuming three neutrino flavours, in figure 2 (the $^4$He scale has been expanded by a factor of 10 for convenience). The effect of increasing theoretical lithium production by a factor of 3 (Olive *et al.* 1981) is also shown. With the error estimates I have used, all the abundance data are consistent with a present-day baryon density (assuming $T = 2.7$ K) that is below $2.3\times 10^{-31}$ g cm$^{-3}$, this limit being set by $Y_p(^4\text{He}) < 0.25$. Lower limits to the density come from the maximum primordial abundances of D, $^3$He and $^7$Li, all of which involve rather more uncertainties, but suggest a limit somewhere near $10^{-31}$ g cm$^{-3}$. If the helium abundance actually turns out to be $Y_p = 0.22$, the upper limit to the density is only $5 \times 10^{-32}$ and this would then be inconsistent with the upper limit adopted for deuterium. Should this happen, the deuterium limit may need to be re-appraised; the primordial deuterium abundance would then be as high as $10^{-3}$ in the canonical model, and this should be relatively easy to detect if it is actually present in the quasar Lyman systems.

TABLE 4. PRESENT-DAY AND PRIMORDIAL ABUNDANCES (BY MASS)

| | $X$ | $X_p$ | | |
|---|---|---|---|---|
| | (now or proto-solar) | lower limit | preferred value | upper limit |
| hydrogen, $^1$H | 0.70 | 0.75 | 0.77 | 0.79 |
| deuterium, $^2$H | $3 \times 10^{-5}$ | $2 \times 10^{-5}$ | $5 \times 10^{-5}$ | $2 \times 10^{-4}$ |
| helium, $^3$He | $6 \times 10^{-5}$ | 0 | ? | $9 \times 10^{-5}$†<br>$4 \times 10^{-4}$‡ |
| helium, $^4$He | 0.28 | 0.21 | 0.23 | 0.25 |
| lithium, $^7$Li | $5 \times 10^{-9}$ | $2 \times 10^{-10}$ | $5 \times 10^{-10}$ | $5 \times 10^{-9}$ |

† If there is no net destruction in stars.
‡ If stars destroy $^3$He completely.

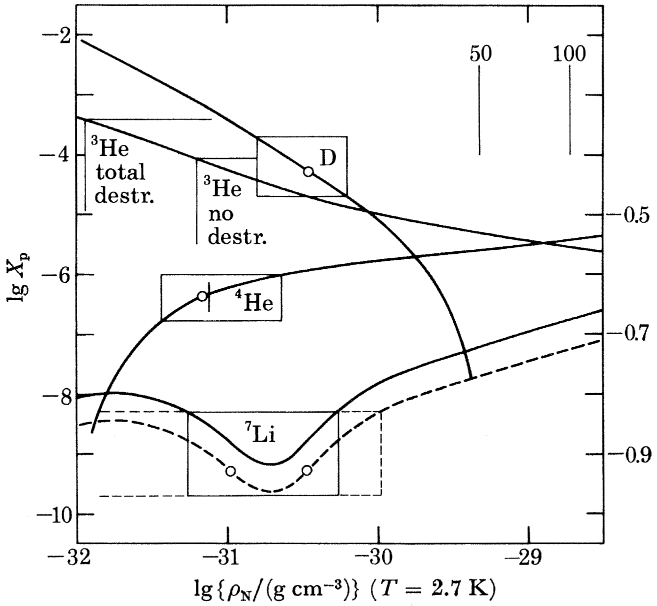

FIGURE 2. Comparison of primordial abundance estimates of this paper with the predictions of the canonical three-neutrino Big-Bang model (Yang *et al.* 1979). The logarithm of $X_p$ is plotted against the logarithm of the average baryon density when the microwave background temperature is 2.7 K. Solid and broken lines for $^7$Li indicate the effect of an upward revision by a factor of 3 (Olive *et al.* 1981). The $^4$He scale is expanded by a factor of 10 for convenience. Open circles indicate preferred values and the corresponding densities, with error boxes according to table 4. Vertical bars in the top right corner indicate closure densities if the Hubble constant is 50 or 100 km s$^{-1}$ respectively, while the vertical mark on the He curve indicates the effect of one neutrino flavour more or less.

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## *Discussion*

G. R. ISAAK (*Department of Physics, University of Birmingham, U.K.*). Our value (Claverie *et al.* 1981) for the spacing between successive $l = 0$, $l = 1$ or $l = 2$ modes in the acoustic spectrum of global solar oscillations with periods near 5 min is 135.2 μHz, and not 136.0 μHz as used by Christensen-Dalsgaard & Gough (1980) on the basis of observations of Fossat and collaborators. Our value has, in the meantime, been confirmed by Hudson & Woodard (1982) using data from solar irradiance studies on the Solar Maximum Mission. Our spacing, interpreted in terms of the standard solar model of Iben & Mahaffy (1976), gives a very low ($Z < 0.005$) heavy element and helium abundance ($Y < 0.16$). Similar results follow from the dirty solar model of Christensen-Dalsgaard *et al.* (1979). A modified version of the latter model (1980) still gives a $Y = 0.19$ for a spacing of 135.2 μHz.

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