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BIG BANG NUCLEOSYNTHESIS AND ABUNDANCES OF LIGHT ELEMENTS

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Abstract. Big Bang nucleosynthesis (BBNS) theory is sketched, indicating the dependence of primordial abundances of D, ^{3}He, ^{4}He and ^{7}Li on the mean baryonic density of the universe and the dependence of ^{4}He on the number of neutrino families and the neutron half-life. Observational data and inferred primordial abundances of these elements are reviewed and shown to be consistent (within errors) either with standard BBNS in a homogeneous universe about 100 seconds after the Big Bang or with moderately inhomogeneous BBNS models resulting from earlier phase transitions like the quark-hadron transition if this is first order. However, models with closure density supplied by baryons are apparently ruled out. Finally, implications for the existence of baryonic and non-baryonic dark matter are briefly discussed.

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1. INTRODUCTION

1.1. The hot Big Bang

Observations of the Hubble expansion and the very isotropic microwave background suggest that the universe has evolved from an earlier state of high temperature and density (*hot Big Bang*) that can be reasonably well described by Friedman-Lemaitre-Robertson-Walker cosmological models. Further support for this view comes from observations of the light elements D, ^{3}He, ^{4}He and ^{7}Li, which are expected to have been synthesised in significant quantities in nuclear reactions that set in about 100 seconds after the Big Bang, an effect that was first suggested by Gamow and his collaborators in the late 1940's although their aim of explaining all elements in this way could not be realised. The modern theory has been developed by Peebles (1966), Wagoner, Fowler & Hoyle (1967) and Yang et al. (1984), among others, and is described by Schramm & Wagoner (1979), Tayler (1982), Boesgaard & Steigman (1985) and Kolb & Turner (1990).

Since the mass density of radiation and relativistic particles varies with scale factor R or red-shift z as $(1 + z)^4$ and that of non-relativistic matter only as $(1 + z)^3$, the gravitating matter of the universe was dominated by the former during the first 10^5 years or so and the universe was then in a phase with significant pressure but negligible effects from curvature or the cosmological constant (if any), and the total mass density at any one time (essentially all radiation and relativistic particles) is then fixed:

$$\rho = \frac{3}{32\pi G} t^{-2}. \qquad (1)$$

The density, in turn, fixes the radiation temperature $T\gamma$ through the equation of state which depends on the number of relativistic degrees of freedom thermal equilibrium with photons. About ls ABB, with a temperature of th. order of 1 Mev, we have, in comparable numbers, photons, electrons, positron and N_ν kinds of pairs of neutrinos and antineutrinos, all of which are relativistic, and a small sprinkling of non-relativistic protons and neutrons, leading t the equation of state

$$\rho c^2 = \frac{\pi^2}{30} \frac{(kT_\gamma)^4}{(\hbar c)^3} \left[g(bosons) + \frac{7}{8} \sum g_i \left(\frac{T_i}{T_\gamma} \right)^4 (fermions) \right] \qquad (2)$$

$$= aT_\gamma^4 [1 + \frac{7}{4} + \frac{7}{8} N_\nu (T_\nu / T_\gamma)^4], \qquad (3)$$

where g, g_i represent statistical weight factors and a is the usual Stefan Boltzmann radiation density constant. The *one* in equation (3) comes from photons, the 7/4 from electrons and positrons and the third term from th neutrinos (and any other, hypothetical particles that might be relativistic at temperature of a few Mev and would then act like an additional contribution to N_ν). With $N_\nu = 3$ and $T_\nu = T\gamma$, (1) and (3) lead to the temperature law (with t in seconds)

$$T = T_\gamma = T_\nu = 0.96 \times 10^{10} t^{-1/2} \text{ K} \qquad (4)$$

$$kT = 0.8 t^{-1/2} \text{ Mev}$$

After several seconds, neutrinos have decoupled and electrons and positrons annihilate, adding entropy to the photon gas, whereafter $T\gamma^3$ is a factor 11/4 greater than $T\gamma^3$. After annihilation, nucleons and photons are both conserved in a co-moving volume so that their ratio $\eta = n_b / n\gamma$ has remained constant (a few times 10^{-10}) through the epoch of nucleosynthesis and to this day. Since the radiation temperature and total density are fixed functions of time, the outcome of primordial nuclear reactions (depending on particle densities and velocities as functions of time) can be expressed as a function of the one cosmic parameter η, apart from physical constants which (in principle at least) can be directly measured in the laboratory.

1.2. Primordial Nucleosynthesis

Nuclear reactions set in when the temperature is down to about 0.1 Mev, since before that deuterium is prevented by photo-disintegration from building up to a sufficient abundance to allow further reactions to occur (the *deuterium bottleneck*). Physical constants affecting the outcome are, naturally, the relevant nuclear reaction cross-sections and two factors affecting the neutron-proton ratio at the onset of synthesis, namely N_ν and the weak interaction constant G_F; G_F^2 is inversely proportional to the half-life of the neutron, $\tau_{1/2}$, which has to be found from experiment.

Fig. 1 shows the predicted abundances from primordial nucleosynthesis according to Standard (i.e. homogeneous) Big Bang Nucleosynthesis (SBBN) theory as a function of η, or, equivalently, of $\Omega_{b0} h_0^2$ where Ω_{b0} is the fraction of the cosmological closure density $3H_0^2 / (8\pi G) = 1.96 \times 10^{-29} h_0^2$ gm cm^{-3} supplied by baryons, H_0 is the Hubble constant and h_0 or h_{100} the same in units of 100 km s^{-1} Mpc^{-1} which lies somewhere between 0.5 and 1. η is related to Ω_{b0} through the known temperature of the microwave background

$$\Omega_{b0} h_0^2 = 3.73 \times 10^{-3} (T_{\gamma 0} / 2.75 \text{ K})^3 \eta_{10}, \qquad (5)$$

where the suffix zero refers to the present epoch and η_{10} is η in units of 10^{-10}. The trends in fig. 1 arise from the series of nuclear reactions starting with

$$\begin{aligned} p + n &\rightarrow D + \gamma, \\ D + D &\rightarrow {}^3He + n \text{ and } T + p, \\ D + (p, T, {}^3He) &\rightarrow {}^3He + \gamma, {}^4He + n, {}^4He + p), \\ {}^3He + {}^3He &\rightarrow {}^4He + 2p, \end{aligned}$$

where the nuclear chain is temporarily halted because there are no stable nuclei with mass numbers 5 and 8. The main effect is to build up ^{4}He with a mass fraction

$$Y_p = 2n/(n + p) \simeq 0.25, \qquad (6)$$

where n / p is the neutron-proton ratio at the onset of synthesis, but traces of D and ^{3}He survive because expansion and cooling slow down nuclear reactions before their destruction is complete. The two-body character of the reactions that destroy deuterium leads to the steep decrease in its abundance with η. Later traces of ^{7}Li and ^{7}Be are also built up, the latter eventually decaying to ^{7}Li by *K*-capture:

$$\begin{aligned} {}^4He + T &\rightarrow {}^7Li + \gamma, \\ {}^4He + {}^3He &\rightarrow {}^7Be + \gamma, \\ {}^7Be + n &\rightarrow {}^7Li + p, \\ {}^7Li + p &\rightarrow 2{}^4He. \end{aligned}$$

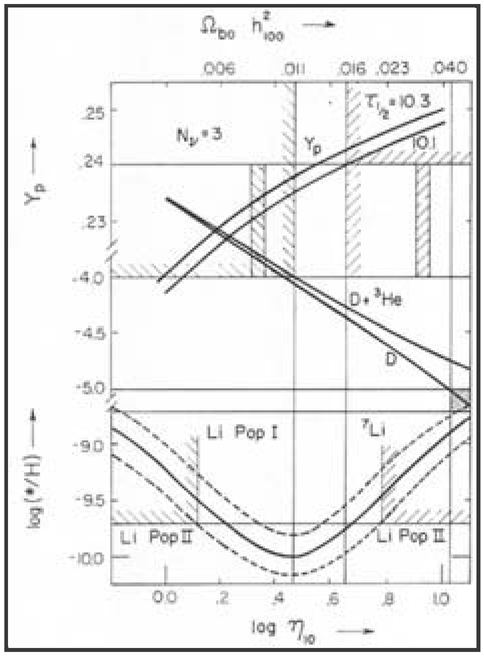

Figure 1. Primordial abundances predicted from SBBN theory after Yang et al. (1984). Olive et al. (1990) and Deliyannis et al. (1990), ^{7}Li from the latter reference being shown with ±2 σ error limits, as functions of η and $\Omega_{b0}\, h_{100}^{2}$. Helium abundances are given for two possible values of the neutron half-life assuming N_{ν} = 3. Horizontal lines show upper limits (and a lower limit for D) based on observation and on reasonable galactic chemical evolution considerations. Tal vertical lines show the corresponding limits on the density parameters from SBBN while the shorter double vertical lines show approximate limits from midly inhomogeneous models adapted from Kurki-Suonio et al. (1990).

At low densities, the first of these reactions is the main contributor to ^{7}Li and its abundance is a decreasing function of nucleon density because *T* and ^{7}Li are both destroyed in two-body reactions. At higher densities the second reaction predominates; ^{3}He and ^{7}Be are more robust, so that the ^{7}Li curve turns round and rises at higher densities leading to a minimum in the interesting range.

Helium itself increases only very slowly with η because virtually all the neutrons initially present are soaked up in its production, as expressed by equation (6). However, its abundance is also significantly affected by N_{ν} and $\tau_{1/2}$ because of the effects of these constants on the initial *n* / *p* ratio. Before (electron) neutrino decoupling, this ratio is kept in thermal equilibrium by weak interactions so that

$$n/p = e^{-1.3\ \mathrm{Mev}/kT}, \qquad (7)$$

whereas afterwards the ratio is virtually frozen at the value given by (7) with *T* equal to the decoupling temperature T_{d}. Because the weak-interaction cross section is proportional to (energy)^{2} or T^{2}, the reaction time scale varies as $(n_{\nu}\, G_{F}^{2}\, T^{2})^{-1}$ or $\tau_{1/2}\, T^{-5}$, where n_{ν} is the number density of electron neutrinos, whereas the expansion time scale is ~ $(G\, \rho)^{-1/2}$ proportional to $(11/4 + (7/8)N_{\nu})^{-1/2}\, T^{-2}$, so the decoupling temperature varies as $\tau_{1/2}^{1/3}(11/4 + (7/8)N_{\nu})^{1/6}$. This means that larger values of $\tau_{1/2}$ or N_{ν}, lead to higher T_{d} and hence to larger *n* / *p* and larger primordial helium abundances. After decoupling, *n* / *p* decreases only slowly through free decay (and residual weak interactions) leading to a very slow dependence on η given by

$$Y_p = 0.228 + 0.023 \log \eta_{10} + 0.012(N_\nu - 3) + 0.018(\tau_{1/2} - 10.28), \qquad (8)$$

(Olive et al. 1990) for 2.5 $\leq \eta_{10} \leq$ 10 and with $\tau_{1/2}$ in minutes. Recent experimental values of $\tau_{1/2}$ are 10.25 (Mampe et al 1989; Gudkov et al. 1990) and 10.32 (Byrne et al. 1990). The sensitivity to N_ν, first noted by Hoyle & Tayler (1964), and the relatively high degree of precision to which Y_p can be estimated (see below), have enabled cosmological limits to be placed on the number of neutrino families, first no more than 5 (Steigman, Schramm & Gunn 1977), then 4 (Yang et al. 1984) and finally 3 (Pagel 1988; Pagel & Simonson 1989; Olive et al. 1990) which was confirmed in accelerator experiments on the Z^o (e.g. Ellis, Salati & Shaver 1990), but with the provisos that SBBN does not exclude neutrinos massive enough to be non-relativistic at a few Mev, which are excluded up to 45 Gev by Z^o decay, and conversely SBBN does exclude hypothetical light particles coupling to photons but not to the Z^o. Consistency of SBBN theory also imposes an upper limit on $\tau_{1/2}$ of 10.4 minutes. Upper and lower limits to η and $\Omega_{b0}\, h_0^2$ resulting from a comparison of SBBN predictions with primordial abundances deduced by various means from observations, as discussed below, are shown by the tall vertical lines in fig. 1.

1.3. Non-standard BBNS models

In recent years there has been active discussion of alternative, non-standard BBNS models that postulate baryon density fluctuations arising from the quark-hadron phase transition (if this is indeed first-order) and related variations in the *n* / *p* ratio due to differential diffusion of protons and neutrons (Applegate & Hogan 1985; Alcock, Fuller & Mathews 1987; Applegate, Hogan & Scherrer 1988; Kawano, Fowler & Malaney 1990). In particular, it has been suggested that such models could fit light-element abundances with Ω_b = 1. The analysis involves a number of free parameters (density contrast, filling factors and length scales) and the proper treatment of all diffusion effects is difficult and controversial, but at the present time models with Ω_b = 1 do not seem to be viable because they predict too much helium and lithium 7 for any combination of the free parameters (Terasawa & Sato 1989, 1990; Reeves 1988, 1990; Kurki-Suonio et al. 1990). However, Reeves (1990) and Kurki-Suonio et al. find that mildly inhomogeneous models are quite plausible and could fit the data for somewhat wider limits on η than are given by SBBN; a rough adaptation of these wider limits from the latter reference is shown by the shorter double vertical lines in fig. 1. Upper limits on N_ν, and $\tau_{1/2}$ are affected very little in these models. A remote possibility exists that there might be significant primordial abundances of elements above ^{7}Li from some kind of inhomogeneous BBNS, but existing data certainly do not suggest anything of the sort (Pagel 1991).

A completely different type of non-standard BBNS theory involves hypothetical massive, unstable particles (e.g. photinos, massive neutrinos, antimatter etc.) which could have various effects depending on their mass, interaction strength and lifetime. For example, they could modify the equation of state, and the success of SBBN and accelerator experiments now rule out large regions of parameter space. They might also decay before, during or after BBNS, modifying the final products. Dimopoulos et al. (1988) suggested that massive (> 2 Gev) particles decaying after 10^5 s (early enough not to disturb the microwave background) produce electromagnetic and hadron showers which *wipe the slate clean* after BBNS and remove SBBN restrictions on Ω_b and N_ν. This particular model makes detailed predictions that disagree with astrophysical observations, in particular too high a ratio of ^{6}Li to ^{7}Li (Audouze & Silk 1989), but these models are generically unappealing on the more fundamental grounds that they perversely throw away the impressive predictions of SBBN theory.

1.4. Primordial abundances

The determination of primordial abundances, which will be discussed in some detail below, requires solution of two separate problems:

1. Determination of abundances in objects that can be sampled or observed now; and **2.** Extrapolation of these abundances back to pregalactic values that can be identified with products of BBNS, allowing for enhancement or destruction by astrophysical processes that may have taken place in the meantime. In what follows, I consider the separate elements and describe what has been deduced, updating previous discussions of this topic by Pagel (1982), Shaver, Kunth & Kjar (1983), Yang et al. (1984), Boesgaard & Steigman (1985) and Pagel (1987a).

2. DEUTERIUM AND ^{3}He

2.1. Deuterium

Deuterium is particularly interesting because stellar processing and the recycling of gas through stars (*astration*) generally cause it to be destroyed and not created, and its very existence is an argument for the Hot Big Bang as was originally stressed by Gamow at a time when its presence had only been established in terrestrial and meteoritic water where it is enhanced by a factor of about 6 owing to fractionation. That this is so was established from studies of the solar wind in meteorites and lunar foils and soils from 1970 onwards (Black 1971, 1972; Geiss & Reeves 1972; Boesgaard & Steigman 1985). The Solar wind contains ^{3}He inherited from the interstellar medium (ISM) when the Solar System was formed 4.6 Gyr ago; ^{3}He resulting from destruction of proto-solar deuterium; and (perhaps) ^{3}He dredged up by turbulent mixing from nuclear-processed material deep inside the Sun (Schatzman 1987). The result is ^{3}He / ^{4}He = (4.0 ± 0.2 (s.e.)) × 10^{-4} or, assuming He/H = 0.1 in the Sun, $y_{23} \equiv {}^3$He / H = (4.0 ± 0.2) × 10^{-5}. On the other hand, gas released by heating carbonaceous chondrites, believed to represent solar wind particles implanted near the time of the birth of the Solar System, contains a smaller proportion of ^{3}He, ^{3}He / ^{4}He = (1.52 ± .05) × 10^{-4}, corresponding to the proto-solar abundance y_3 of ^{3}He on its own. If fresh production of ^{3}He by dredge-up is ignored, the proto-solar D / H ratio is simply the difference y_{23} - y_3 = (2.5±0.2) × 10^{-5}, with which ground-based and Voyager infrared observations of deuterated molecules (HD, CH_3D) in the atmospheres of Jupiter, Saturn and Uranus, carried out since 1972, are in fair agreement (see Boesgaard & Steigman 1985; Pagel 1987a). Interstellar deuterium was discovered in 1973, in molecular form from radio observations of molecular clouds (Jefferts, Penzias & Wilson 1973) and as HD and DI (Lyman bands and Lyman series) from ultra-violet spectroscopy of diffuse clouds in front of hot stars using the Copernicus satellite (Spitzer et al. 1973; Rogerson & York 1973). A relatively low abundance of DCO^+ and DCN at the Galactic centre (Penzias 1979) supports the purely (or at least mainly) destructive effect of astration on deuterium, but unknown fractionation effects make it difficult to infer the interstellar D / H ratio from molecules except in the case of DCO^+ (Dalgarno & Lepp 1984) which agrees with atomic lines in giving a ratio ~ 10^{-5}. A hyperfine transition of DI, at 91.6 cm wavelength, has been searched for several times, but without a definite detection (Pasachoff & Vidal-Madjar 1989). Most determinations of the interstellar D/H ratio come from observations of Ly γ, δ, ϵ in absorption on lines of sight to hot stars at distances up to 1 kpc and a few from IUE observations of Ly α emission lines, with interstellar absorption superposed, from very nearby stars, the deuterium line appearing as a weak component displaced to the violet by 81 km s^{-1}. Difficulties arise from appropriate modelling of the velocities and velocity dispersions of the intervening clouds (done with the help of optical observations of *NaI*) and from the possibility of spurious signals arising from hydrogen clouds expelled at about 80 km s^{-1} from the target star, for which there is direct evidence in some cases (Vidal-Madjar et al. 1983; Gry, Lamers & Vidal-Madjar 1984). Estimates of the interstellar D / H ratio thus cover quite a wide range, from 2.5 × 10^{-5}, the same as the proto-solar value, to 6 × 10^{-6} (Boesgaard & Steigman 1985; Pasachoff & Vidal-Madjar 1989).

2.2. Helium 3

^{3}He has been detected in a solar prominence from its isotope shift relative to the 10830 emission line with an estimated abundance consistent with that in the Solar wind (Hall 1975). ^{3}He is also observed in the ISM, in HII regions, using large radio dishes, by virtue of the hyperfine transition of $^3He^+$ at 3.46 cm (Bania, Rood & Wilson 1987). Abundance determinations relative to nearby H recombination lines involve correction for clumping and for differing degrees of ionisation of hydrogen and helium, but the most difficult part is subtraction of the baseline from the data, since the line is exceedingly weak. The resulting abundance ratios ^{3}He / H range from 1.5 × 10^{-4} to under 10^{-5}, without any clear pattern with galactocentric distance such as might have been expected from the trends of oxygen and nitrogen abundance. Some (though not necessarily all) of the source-to-source variations appear to be real.

2.3. Primordial D+ 3He

How do these results on D and ^{3}He relate to primordial abundances? Since D is mainly destroyed by astration, its proto-solar abundance can be taken as a lower limit, but estimates of how much lower than primordial are rather model-dependent with respect to Galactic chemical evolution, depending on how much of the original Galaxy is still in the form of gas (10 to 20 per cent in the Solar neighbourhood), how much gas is returned to the ISM by each generation of stars (10 to 40 per cent) and whether there has been significant inflow of unprocessed gas from outside (cf. Pagel 1982). Thus factors from 2 to 10 are all quite conceivable. In the case of ^{3}He, the situation is still more difficult because ^{3}He is destroyed in astration through massive stars, but survives and can also be freshly produced in stars of lower mass (Dearborn, Schramm & Steigman 1986). Yang et al. (1984) pointed out that, because destruction of D leads to production of ^{3}He, some of which survives further stellar processing, one can use existing abundances in the proto-Solar System to constrain the sum of primordial D and ^{3}He by the equation

$$y_{23p} \leq y_{23p} + (\frac{1}{f} - 1)y_3 \quad (9)$$

$$\leq (5.5 \pm 0.2) \times 10^{-5} \text{ if } f = 0.5 \text{ (likely)} \quad (10)$$

$$\leq (8.6 \pm 0.3) \times 10^{-5} \text{ if } f \geq 0.25 \text{ (almost certain)} \quad (11)$$

where y_{23p} is the sum of primordial (D + ^{3}He) / H and f is the fraction of ^{3}He that survives astration through one generation of stars. In a somewhat more sophisticated treatment the second limit is increased to 10.9 × 10^{-5} (Olive et al. 1990), which is shown to a good approximation by a horizontal line in fig. 1. This gives what is currently the most stringent lower limit to η in the framework of BBNS, shown by the left-most tall vertical line and the leftmost shorter double vertical line in fig. 1 for homogeneous and inhomogeneous models respectively. Equation (9) refers to closed Galactic chemical evolution models without inflow of unprocessed material. In models with inflow, which have some distinct advantages (Pagel 1989a), different arguments apply but the result is the same, since in such models it is virtually impossible to have less than 1/3 of primordial deuterium surviving (cf. Audouze & Tinsley 1974; Pagel 1982).

It would be nice to have an independent upper limit to primordial deuterium and there is a possibility that this will eventually be achieved from observations of absorption-line systems at high red-shifts having low metallicity, low Doppler broadening and large hydrogen column density, in front of quasars (Webb et al. 1991). The absence of any very definite results on deuterium in such clouds up to now suggests that the limit of 10^{-4} is not too likely to be violated.

3. LITHIUM 7

3.1. Lithium in the galaxy

^{7}Li is readily detected spectroscopically from the resonance doublet of LiI at 6707.8 Å and has been found in hundreds of stars cooler than effective temperature 7500 K (spectral type F0) and in the ISM as well as in meteorites where its abundance is equivalent to Li/H = 2 × 10^{-9}. A similar abundance is found in the photospheres of young stars (if hotter than about 5500 K, spectral type G6), but in older stars lithium is depleted due to mixing or dilution with deeper layers where ^{7}Li is destroyed by the (p, α) reaction at temperatures above 2.5 × 10^6 K. Because cooler stars have deeper subphotospheric convection zones, they deplete their surface lithium faster so that, in the young Pleiades cluster (age ~ 10^8 yrs), Li has its normal abundance down to effective temperature 5500 K, whereas in the older Hyades cluster (~ 10^9 yrs) depletion sets in and intensifies steadily below 6300 K (type F7). In the Hyades and NGC 752, furthermore, there is also a high degree of depletion in a narrow range of effective temperatures around 6600 K (Boesgaard & Trippico 1986; Boesgaard & Budge 1988; Hobbs & Pilachowski 1988) near the region of pulsational instability, possibly caused by mass loss; the amount of mass actually lost has to be quite small because beryllium is not significantly depleted in these stars (Schramm, Steigman & Dearborn 1990). In cooler stars, there is evidence that the depletion depends on other factors besides mass and age, since the degree of depletion in the solar photosphere, which is a factor 100 or so, is greater than in some stars of similar effective temperature in the still older galactic cluster NGC 188 (Hobbs & Pilachowski 1988; Hobbs, Iben & Pilachowski 1989); such factors may include rotationally driven meridional circulation, magnetic effects etc. Thus ^{7}Li is subject to destruction by astration, but it is certainly synthesised in a subset of carbon stars where it is observed to be superabundant (Smith & Lambert 1991) and possibly in novae and there is a minor contribution due to cosmic-ray spallation; the net effect of all these processes is quite likely to lead to an increase in ^{7}Li abundance in the ISM as a result of stellar activity (see fig. 2).

3.2. Lithium in subdwarfs

The big breakthrough in relating ^{7}Li to cosmology came with the brilliant discovery by Spite & Spite (1982a, b) of Li in metal-deficient subdwarf stars of the Galactic halo (extreme Population II). Relative to the Sun and most nearby stars, these are deficient in carbon and heavier elements by factors between 10 and 3 × 10^4 and they are the oldest stars that we know, and yet in the range of effective temperatures 5500 K to 6500 K (where the main sequence turns off), as long as the metal-deficiency factor exceeds 50, they have virtually constant lithium abundance $-10.2 < \log(Li/H) < -9.7$ only a factor of 10 below young Population I stars (Rebolo, Molaro & Beckman 1988; cf. fig. 2). Below 5500 K, the Li abundance goes down with diminishing effective temperature, analogously to what is found in Pop. I.

Since the subdwarfs consist of virtually pristine material, it is natural to follow Spite & Spite in assuming that their lithium is all primordial ^{7}Li (it is mainly ^{7}Li and not ^{6}Li; Spite, Maillard & Spite 1984) and the only question that arises is whether it represents the full primordial abundance or whether after such a long time since the formation of these stars the photospheric abundance has been depleted. The fact that the abundance has a constant *plateau* over a wide range in effective temperature suggests that the amount of depletion should be small, and this is indeed what is predicted by standard stellar evolution calculations (Deliyannis, Demarque & Kawaler 1990) which include diffusion but no non-classical effects like meridional circulation some of which must have been operating in Pop. I stars such as the Sun. Thus significant depletion is not entirely ruled out (Vauclair 1988; Mathews, Alcock & Fuller 1990; Krauss & Romanelli 1990), but it could very well be absent and, if so, then subdwarf Li gives a very impressive fit to SBBN predictions (which preceded its discovery) and limits η from both above and below 1).

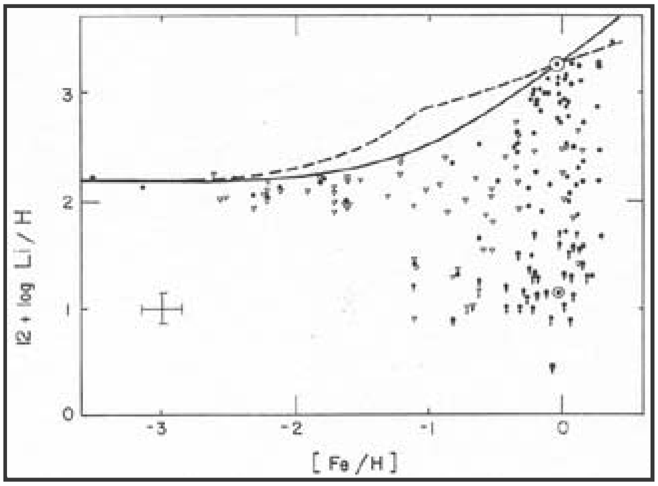

Figure 2. Plot of stellar ^{7}Li abundances after Rebolo, Molaro & Beckman (1988). solid curve shows the prediction of a simple model in which ^{7}Li is the sum of a primordial component and an additional component proportional to iron. The broken curve assumes the additional component proportional to oxygen. The two Sun symbols refer to meteoritic (identified as proto-solar) and photospheric (depleted) values respectively.

4. HELIUM 4

4.1. Introduction

Helium is the second most abundant element in the visible universe and there is accordingly a vast amount of information about its distribution from optical and radio emission lines in nebulae, optical absorption lines in spectra of hot stars (effective temperatures above 10^4 K), scale heights of the atmospheres of the major planets and the influence of initial helium content on stellar structure and evolution and pulsation. However, primordial helium has been enhanced by astration to varying degrees in different objects, so that the observed mass fraction Y is in general only an upper limit to the primordial mass fraction Y_p, and, conversely, Y may be reduced in planetary and certain stellar atmospheres by gravitational settling. Thus in the 1960's a certain amount of confusion reigned as to whether there was really a universal lower limit to Y as required by SBBN (cf. Pagel 1982). The existence of hot subdwarfs with weak helium lines was one source of confusion until Sargent & Searle (1967) showed that other abundance anomalies in these stars implied processes similar to those operating in the chemically peculiar A-type stars of Population I. Christy (1966) had already shown that the pulsational characteristics of RR Lyrae stars, many of which occur in halo globular clusters, demand a substantial helium content in the envelope and Iben (1968) likewise deduced a substantial initial helium content from the relative numbers (i.e. lifetimes), in globular clusters, of core helium-burning stars on the horizontal branch of the luminosity-effective temperature diagram (the Hertzsprung-Russell diagram) and hydrogen shell-burning stars on the red giant branch. Searle & Sargent (1972) finally clinched the matter by taking spectra of two blue compact galaxies discovered by Zwicky, I Zw 18 and II Zw 40, and showing that their light is dominated by H II regions with very low abundances of observable heavy elements (N, O, Ne and S), but nearly normal helium.

A thorough account of all the data available up to 1983 has been given in the proceedings of the ESO Workshop (Shaver, Kunth & Kjar 1983) and table 1 gives a slightly updated version of the results presented there, which strongly support the view that there is a universal floor corresponding to a primordial abundance $0.20 < Y_p < 0.25$, but for reasons briefly indicated in the last column of the table it is difficult to reach the better than 5 per cent precision (i.e. less than ±.01) that is needed in order to constrain the SBBN model significantly. Fig. 1 shows that, with the lower limit to η derived from D + ^{3}He, and standard values of N_ν and $\tau_{1/2}$, Y_p needs to be at least 0.235 for consistency with SBBN and that a firm upper limit to Y_p may give a tighter upper limit to η than can be derived from other elements. However, the Sun and hot stars give only a rather imprecise upper limit and in subdwarfs and halo globular clusters, which are likely to approximate pristine material quite closely, the relevant stars are too cool to show helium lines in their spectra and one has to rely on systematics of the HR diagram: location of the zero-age main sequence (ZAMS, needing knowledge of distance); effective temperature at the blue edge of the instability strip where RR Lyrae variables are found; the difference in magnitude between the horizontal branch (HB) and a well-defined point on the ZAMS; and the relative numbers of stars in the HB and red giant stages. These estimates are dependent on a full understanding of stellar opacities, reaction rates, instability mechanisms, convection, semi-convection etc. (cf. Caputo 1985), so that the systematic uncertainties are difficult to judge.

Table 1. Some estimates of (or upper limits to) primordial helium

Obj.observed	Y_p	Method	First author	Problems
Sun	< .28±.02	Struc.; oscill.	Turck-Chieze 88	Eq. of state
	< .28±.05	Promiinence em.lines	Heasley 78	Self-abs.
Hot stars	< .28±.04	Absorption lines	Wolff 85	Precision
Subdwarf	.19±.05	Lum. and eff. temp	Carney 83	Parallax; convection
Glob.clust.	.23:	Variables and lum.evol.	Caputo 83	Physical basis
	.20±.03 :	N(HB)/N(RG)	Cole 83	of stellar structure
	.23±.02	" : "	Buzzoni 83	and evolution
	.24±01	" : "	Caputo 87	(see Caputo 1985)
Plan.nebula	.22±.02	opt.em.lines	Peimbert 83	Self-and Gal. enrichment
Gal. HII reg	.22:	opt.and radio em.lines	Mezger 83	He^0 and Gal. enrichment
Extragal.	.233±.005	opt. em.lines	Lequeux 79	(See text)
H II region	< .243±.010	" : "	Kunth 83	

4.2. Recombination lines in nebulae

More precise estimates of helium abundance come from observations of emission lines in gaseous nebulae (planetary nebulae and H II regions), where both hydrogen and helium lines are formed predominantly by recombination of H^+ and He^+ for which a precise theory exists (Brocklehurst 1972) and adequate signal: noise can be obtained for λ 4472 (4^3D - 2^3P°), λ 5876 (3^3D - 2^3P°) and λ 6678 (3^3D - 2^3P°) using panoramic linear detectors (photon counters or CCD's) on nebulae with sufficiently high surface brightness. For Galactic H II regions, radio recombination lines can also be used (Thum, Mezger & Penkonin 1980; Thum 1981; Peimbert et al. 1988). However, Galactic H II regions have rather large heavy-element abundances, which makes extrapolation to pregalactic values uncertain, and planetary nebulae are additionally affected by the internal evolution of their central stars. Thus the most favourable objects for the estimation of Y_p are extragalactic H II regions in dwarf galaxies (or the outer parts of spirals) where the heavy-element abundances are low. Bright examples of this class are seen in a few spirals (e.g. M101), in nearby irregular galaxies like the Magellanic Clouds, in a subset of blue compact galaxies discovered on direct photographs by Zwicky and Haro, and in H II galaxies, dominated by emission lines, and mostly discovered in objective prism surveys in Armenia (Markarian) and Chile (Palomar, Michigan and Cerro Tololo surveys).

Peimbert & Torres-Peimbert (1974, 1976) noticed a small but significant trend for helium abundance to increase with heavy-element abundance in the order I Zw 18, SMC and II Zw 40, LMC, Orion Nebula, and accordingly proposed that Y_p could be found by looking at H II regions with different heavy-element abundances, plotting a linear regression of the form

$$Y = Y_p + Z(dY/dZ) = Y_p + (O/H)[dY/d(O/H)], \qquad (12)$$

and extrapolating to $Z = 0$.($Z \simeq 25(O/H)$.) They carried out this programme (Lequeux et al. 1979) and derived $Y_p = 0.23$ and $dY / dZ = 3$, the latter quantity turning out to be rather large compared to expectations from the theory of stellar evolution (Maeder 1984; Serrano 1986). A later survey by Kunth & Sargent (1983) showed no clear evidence for a dY / dZ slope, but this is largely due to the high weight given by them to the 5876 line in II Zw 40, an object heavily reddened by dust in the Milky Way, which is probably affected by absorption due to Galactic NaI (French 1980). The remainder of Kunth & Sargent's data actually show a steep, if ill-defined, slope (cf. Peimbert 1985). Various other data, mostly of inferior quality, quoted, for example, by Boesgaard

& Steigman (cf. Pagel 1989b), led to some doubt as to whether a *dY* / *dZ* correlation actually exists, but since then the situation has been somewhat restored as a result of more careful work by Peimbert, Shields, Terlevich, Pagel and associates (see Pagel 1991).

4.3. Complications in emission-line analysis

Although the hydrogen and helium lines are basically due to a simple recombination process, there are various complications in the precise interpretation of their relative intensities (cf. Davidson & Kinman 1985) even when the non-trivial problems of detector linearity, flux calibration and correction for interstellar reddening have been overcome. These are the following:

1. Unobservable neutral helium in the H^+ region. Because He is abundant enough to soak up its own ionising photons, the degree of ionisation of trace elements like oxygen or sulphur is not a straightforwardly good guide to this effect, which is mainly governed by the effective temperature(s) of the ionising star(s) (Osterbrock 1974) and negligible when this exceeds 40,000 K or so. Vilchez & Pagel (1988), following earlier work by Shields & Searle (1978) and Mathis (1982, 1985), use a *radiation softness parameter* $(O^+ / O^{++}) / (S^+ / S^{++})$ as a measure of the effective temperature and the more recent work avoids cases where the parameter is so large that the corresponding ionisation correction factors (icf) read off from photo-ionisation models (Mathis 1982; Stasinska 1982, 1990) exceed a few per cent and become model-dependent. Because stellar effective temperatures tend to be higher in objects with the lowest heavy-element abundances (e.g. Campbell 1988), there is a danger of a spurious *dY* / *dZ* correlation when objects with a large icf are included. Our method could underestimate the icf in cases where one sees two H II regions superposed, a hot one with HeI, H I and [O III] and a cooler one with H I and [O II] but no, or less, HeI (Pena 1986; Dinerstein & Shields 1986; Dufour, Garnett & Shields 1988), but in our hottest objects [O II] is in any case so weak compared to [O III] that this problem cannot lead to an error as large as 5 per cent (Pagel & Simonson 1989). He^{++} is directly seen by virtue of the λ 4686 line in the hottest objects and is easily allowed for.

2. Collisional contributions to the emission lines. Because the 2^3S state of HeI is highly metastable, it builds up a substantial population in H II regions and can be excited by electron collisions to the upper states of the relevant optical lines (Cox & Daltabuit 1971). This possibility was generally discounted because of agreement between triplets and the singlet λ 6678 in dense planetary nebulae (Peimbert & Torres-Peimbert 1971) until Ferland (1986) drew attention to new quantum-mechanical calculations by Berrington et al. (1985) which implied that the singlet states can also be excited from 2^3S. Ferland deduced some remarkably low helium abundances from λ 5876 after correcting for the effect; these resulted in part from an overestimate of the rates by Berrington et al. (since more exact computation including resonances from enough higher levels is expensive) but in large part also from a poor selection of data from the literature on Ferland's part. Anyway, his results were sensational enough to provoke new, more accurate quantum-mechanical calculations by Berrington & Kingston (1987), giving about half the previous rates, which have been used by Clegg (1987) to provide what seem to be reliable correction formulae depending on electron temperature and density (which therefore need to be accurately measured). Peimbert & Torres-Peimbert (PTP 1987a, b) find that, compared to these formulae, λ 10830 (2^3P^o - 2^3S) is still anomalously weak by a factor 2 or so. Clegg & Harrington (1989) consider various effects that could act to depopulate the 2^3S state, finding that radiative processes do not do so appreciably at moderate or low densities. They accordingly suggest the presence in the planetaries studied by PTP of a hitherto unknown destruction mechanism (like charge exchange) which would be inoperative at the low densities of extragalactic H II regions. The corrections for 4471 and 6678 do not usually exceed a few per cent, in any case.

3. Fluorescence effects. The large population of 2^3S can also lead to enhanced production of emission lines by multiple scattering of ultra-violet photons (Robbins 1970). A line especially sensitive to this is λ 7065 (3^3S - 2^3P^o), which is also relatively strongly affected by collisional excitation, but can still be used as a test when this is allowed for. Existing measurements of λ 7065 (listed by Pagel 1987a) show no evidence for significant fluorescence enhancement of the helium lines from extragalactic H II regions.

4. Underlying absorption lines in the stellar continuum. When an H II region is well resolved, it may be possible to place the slit of the spectrograph in such a way as to record purely nebular emission, in which case this problem does not arise. With H II galaxies, however, which are apparently (and sometimes also intrinsically) very compact, it is impossible to avoid including the spectrum of the embedded star cluster with absorption lines of both hydrogen and helium that are usually unresolved from the nebular emission lines. This problem particularly affects λ 4471 and it can be quantified in high signal:noise spectra by looking for the neighbouring line λ 4388, which is comparable or stronger in absorption-line spectra but only 1/7 as strong in emission. Absorption equivalent widths of hydrogen and helium lines relevant to this problem have been calculated from evolutionary stellar population synthesis models by Olofsson (1990).

5. Effects of internal dust. While extinction effects of external dust along the line of sight are readily allowed for by

comparison of observed and theoretical Balmer decrements using a standard reddening law, internal dust, when present even in small amounts, can cause substantial complications, e.g. by swallowing up Lyman line photons and invalidating the standard assumption of Case B recombination. This problem is quite significant in the case of the Orion Nebula (Cota & Ferland 1988; Baldwin et al. 1991) but probably much less so in extragalactic H II regions with lower densities and lower abundances. The effect, if present, must vanish in the limit $Z = 0$, but it gives yet another reason for not including Orion in the regression (12) (cf. Pagel 1982).

6. Absorption by intervening gas. Both the Milky Way and the Earth's atmosphere have absorption lines that can affect HeI λ 5876 at certain red-shifts, and a particular problem is presented by Galactic NaI for objects with red-shifts in the range 0.002 to 0.004 in which many of the known H II galaxies lie, including both I Zw 18 (Davidson, Kinman & Freedman 1989) and II Zw 40 (discussed above). Because of this, and because of somewhat greater sensitivity to collisional excitation, it is not advisable to try to deduce a helium abundance from λ 5876 alone.

4.4. Newer results

In the last 5 years, I have been actively working on the helium problem, making use of the excellent spectroscopy of H II galaxies by Terlevich & Melnick (Campbell, Terlevich & Melnick 1986; Terlevich et al. 1991), new observational data with emphasis on obtaining high signal:noise in λ 6678 (which is the easiest line to interpret but relatively weak) and what seem to be the best data in the literature, rediscussed in a uniform manner (Simonson 1990; Pagel 1991). Furthermore, we have investigated the correlation with nitrogen as well as oxygen since the correlation with N/H seems to be somewhat better (Pagel, Terlevich & Melnick 1986). Fig. 3 shows one of our best spectra, secured with the Anglo-Australian Telescope in 1988, which gives accurate electron density and S^+ / S^{++} as well as 6678 / Hα and then in combination with blue-yellow spectrophotometry by Terlevich and his associates gives a very secure measurement of the He/H ratio (Pagel & Simonson 1989). So far we have spectra of this sort for only three objects and more are needed.

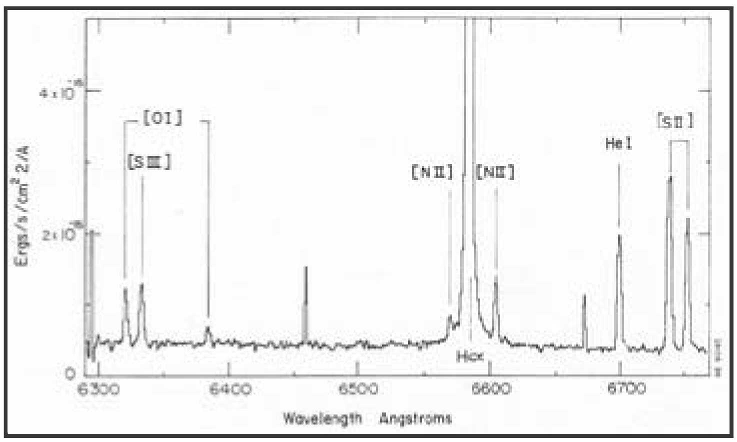

Figure 3. Red spectrum of the H II galaxy UM 461 (Terlevich et al. 1991) taken with the Anglo-Australian. Telescope in April 1988 by Pagel, Simonson & Terlevich, with identifications of emission lines. Narrower spikes are cosmic-ray events in the CCD detector. The spectral resolving power is about 2000.

Fig. 4 shows our regression relations of helium with oxygen and nitrogen in low-abundance extragalactic H II regions with maximum-likelihood linear regression lines and error limits equivalent to $\pm 1\sigma$. The regressions are

$$\begin{array}{rcl} Y = 0.225 & + & 169(O/H) \\ \pm 5 & \pm & 45 \end{array} \qquad (13)$$

and

$$\begin{array}{rcl} Y = 0.229 & + & 3310(N/H) \\ \pm 4 & \pm & 940 \end{array} \qquad (14)$$

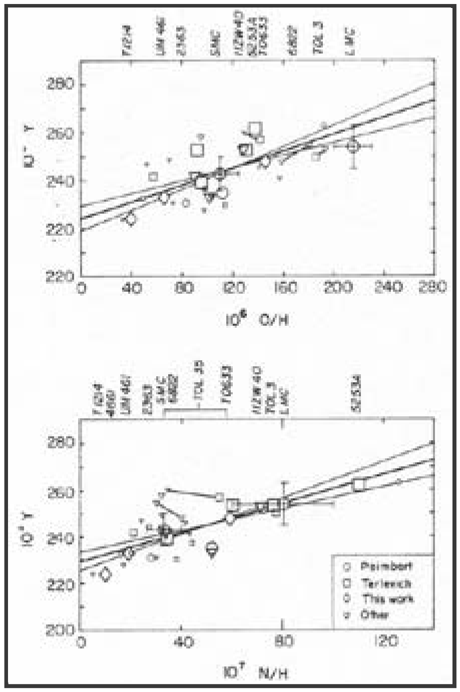

Figure 4. Regressions of helium mass fraction against oxygen and nitrogen abundance, respectively, in irregular and blue compact (or H II) galaxies with oxygen up to 1/4 solar. Maximum-likelihood regression lines are shown with alternatives equivalent to $\pm 1\sigma$ errors. Sizes of plotted symbols indicate their weights; responding error bars typical of hig-weighted data in the same part of the diagram are shown in just two cases. For a tabulation of the data and their sources, see Pagel (1991).

The regression against oxygen has a remarkably steep slope (corresponding to $dY / dZ = 6.5 \pm 2$) and suggestions of either a flattening off towards higher abundances or significant scatter which is absent (or at least not noticeable compared to errors) in the regression with nitrogen (the latter, however, certainly does not continue linearly beyond the range of the diagram; e.g. Orion has $Y \simeq 0.27$, 10^7 N/H $\simeq 700$!). The reason for the better correlation with N/H could be local pollution by winds fron Wolf-Rayet stars in the embedded cluster which produce additional He and N overlying the basic correlation with oxygen noted by Peimbert and his colleagues, whose conclusions we basically confirm. The pollution hypothesis (Pagel, Terlevich & Melnick 1986; Pagel 1987a, b) is supported by a detailed survey of NGC 5253 by Walsh & Roy (1989), but there are still some ambiguities (Simonson 1990). If the effect is not due to pollution, it could perhaps arise from differential galactic enrichment in oxygen, from short-lived massive stars, on the one hand, and in nitrogen and helium, by planetary nebulae from longer-lived intermediate-mass stars, on the other (cf. Edmunds & Pagel 1978; Steigman, Gallagher & Schramm 1990).

4.5. The primordial helium abundance

For whatever reason, the nitrogen regression gives a slightly better correlation than the oxygen one and a slightly more precise result, which is therefore the one that we currently adopt. Table 2 compares it with some previous estimates and it shows that, despite the various arguments raised in Section 4.3, the result has been remarkably stable for over a decade, during which time there have been larger changes in the estimates of the neutron half life (these have come down; see Tayler 1990) and the restriction to $N_{\nu} = 3$ has been confirmed in accelerator experiments. The systematic errors remaining in our estimates of Y_p are inevitably a matter of judgement. I see no reason why they should exceed the formal standard error of 0.004, in which case there is 95 per cent confidence that the true value does not exceed 0.240 as shown by the horizontal line in fig. 1 and Y_p gives the tightest upper limit to the density parameter η. Specifically, assuming SBBN, ((D + ^{3}He/H)$_p \leq 10^{-4}$, $\tau_{1/2} \geq 10.1$ min and $Y_p \leq 0.240$), the limits from equation (8) are

$$2.9 \leq \eta_{10} \leq 4.6, \qquad (15)$$

or, from equation (5),

$$0.011 \leq \Omega_{b0} h_{100}^2 \leq 0.017. \qquad (16)$$

h_{100} is universally agreed to be between 0.4 and 1.0; most probably it exceeds 0.7 (Tully 1990) which rules out an Einstein-de Sitter universe with Ω = 1 because of the ages of globular clusters. If 0.7 $\leq h_{100} \leq$ 1.0 , then

$$0.011 \leq \Omega_{b0} \leq 0.035. \qquad (17)$$

The lower limit calls for baryonic dark matter, since visible matter in spiral and irregular galaxies corresponds to Ω_{vis} = $0.002 h_{100}^{-1}$ and the larger mass:light ratio found in ellipticals is itself probably due in large part to white dwarfs and neutron stars (Yoshii & Arimoto 1987).

Table 2. Estimates of primordial helium abundance with ±1σ errors

.230 ± .004	Lequeux et al. 1979
<.243 ± .010	Kunth & Sargent 1983
.234 ± .008	Kunth & Sargent 1983 without II Zw 40
.232 ± .004	Peimbert 1985
.237 ± .005	Pagel, Terlevich & Melnick 1986
.232 ± .004	Pagel 1987a
.230 ± .006	Torres-Peimbert, Peimbert & Fierro 1989
.229 ± .004	Pagel & Simonson 1989

This amount of dark matter could be present in dark halos of spirals deduced from 21 cm rotation curves and the dynamics of the local group; alternatively, it might be there in the form of low surface-brightness galaxies not counted in conventional optical surveys (Pagel 1990). The upper limit less than most estimates of Ω_0 (total) around 0.2 based on galaxy cluster dynamics (Peebles 1986) leaving some space for non-baryonic dark matter.

In the inhomogeneous BBNS case, the corresponding limits are (rough

$$0.009 \leq \Omega_{b0} h_{100}^2 \leq 0.04, \qquad (18)$$

or with 0.7 $\leq h_{100} \leq$ 1.0

$$0.009 \leq \Omega_{b0} \leq 0.08, \qquad (19)$$

which leaves the case for dark baryornc matter virtually unchanged, but so what weakens that for non-baryonic matter (it would weaken it even more - possibly up to the point of extinction - if h_{100} were smaller!). Whether inhomogeneous case actually applies is unclear; the physical question of extence of a first-order transition still remains to be settled, and then there is question of whether the primordial deuterium and helium abundances and neutron half-life can be squeezed tightly enough to cause real embarrassment to SBBN. Such would be the case, for example, if one could demonstrate exclusively that Y_p < 0.235, but there are enough opportunities for system ore than speculative errors in existing data to make this possibility no more than speculative the time being.

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