Primordial abundances

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Abstract: The Universe's origin in a Big Bang was followed by rapid expansion and cooling. Protons and neutrons precipitated out of the cooling primeval fireball of elementary particles, and were able to fuse together by nucleosynthesis to produce light elements such as deuterium, helium and lithium. The abundances of these elements can be used to constrain models of the Big Bang. The primordial abundances have been modified by stellar nucleosynthesis. Observational determinations of abundances are beset by many problems, and extrapolating these values to a primordial abundance has several complications. The methods used to set limits on primordial abundances are discussed.

1. In the beginning ...

When, where and how did the Universe come into being? The subject has often been incorporated into myths and religion, but the Universe's origin is now a subject of scientific enquiry. With the development of relativity theory and quantum mechanics in the early part of the 20th century, it has become possible to develop explanations for the origin of the Universe in terms of empirically testable models. Coupled to the theoretical developments, there have been technological advances that have made it possible to design and build larger telescopes, with instrumentation that is more sensitive and covers a broader range of the electromagnetic spectrum than was possible in previous generations. Probing the depths of the cosmos has helped to constrain the numerous theories suggested for the origin of the Universe.

Observations made by Edwin Hubble in the 1920s led to his discovery (Hubble 1929) that the further away a galaxy is from the Milky Way, the faster it is receding from us. This velocity-distance relationship, known as the Hubble law, has been confirmed for hundreds of galaxies out to great distances. It shows that the Universe is expanding. Two models of creation were proposed to account for an expanding Universe. The steady state theory, proposed independently by Bondi & Gold (1948) and Hoyle (1948), maintains that the Universe does not evolve and always has the same general properties. To preserve the density of an expanding Universe, the steady state theory required continuous creation of matter. This newly created matter, presumably in the form of hydrogen atoms, collected in great clouds between the receding galaxies and eventually gave birth to new galaxies to take the place of ageing galaxies no longer able to make new stars. In the steady state theory, all elements other than hydrogen are synthesized by nuclear reactions in stars.

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An alternative theory, based on an analysis of an evolving homogeneous and isotropic cosmological model, was put forward by Lemaître in 1927. Similar equations had been derived by Friedmann in 1922, but Lemaître linked his solutions to the observational work being done at Mt Wilson (Lemaître 1931). The basic idea is that if an expanding Universe is extrapolated backwards in time, sometime in the distant past the Universe came into being in an extremely hot, dense state. This primeval fireball had densities and temperatures beyond the limits that current physics can describe, but as the Universe expanded, it cooled. The initial plasma contained elementary particles known as quarks and gluons. After about one second, protons and neutrons were able to condense out of the plasma. Some light elements were produced by nucleosynthesis in the Universe's early phase, while all other elements had to wait for stars to form before they could be synthesized. Sir Fred Hoyle sarcastically proposed the term "Big Bang" for this theory of creation.

Nowadays there is almost unanimous agreement that the Universe came into being between 10–20 Gyrs (Giga = thousand million, or billion) ago in a Big Bang. The rate of expansion and the ultimate fate of the Universe are still subjects of much debate. Was there a period of rapid expansion called inflation in the young Universe, and will the Universe continue to expand forever? These are questions that have not been answered definitively yet. Recent observations have been interpreted as indicating that the rate of expansion is increasing (i.e. accelerating) as the Universe ages.

The theoretical foundation of the Big Bang model consists of using the Robertson-Walker metric in Einstein's field equations to get the Friedmann equations. These equations relate the expansion rate H to the energy density, and the energy density is related to the thermal properties of particles. The standard model of particle physics provides the candidate particles contributing to the energy density. There are a limited number of methods for testing the model, however. Until a few years ago, the three cosmological observables that could be measured most accurately were the Hubble constant H_0, the cosmic microwave background radiation (CMBR) photon density, and the primordial light element abundances. Billion dollar satellite missions have been dedicated to measuring H_0 and the CMBR. H_0 is compatible with the steady state and Big Bang theories, the CMBR and light element abundances are unique to the Big Bang model. The CMBR is a relic of the Universe when it was about 300 000 yrs old, but the light element abundances are a fingerprint from the Universe when it was only 1000 s old, and hence provide the earliest record of the Universe's history.

Whether the Universe will continue to expand forever (an open Universe) or slow down and contract, ending in a big crunch (a closed Universe), is determined by the density of the Universe. The dividing line between an open and closed Universe is a flat Universe. The density of a flat Universe in Friedman-Robertson-Walker models with cosmological constant $\lambda=0$ is called the critical density ρ_{crit} and is related to the Hubble constant by $\rho_{crit} = 3H_0^2 / 8\pi G$. The density of baryons ρ_b relative to the critical density is given by $\Omega_b = \rho_b / \rho_{crit}$.

Elements are distinguished by the number of protons in an atom's nucleus. Protons have a positive charge and repel each other by electrostatic forces. In the nucleus, protons and neutrons are bound together tightly by a force known as the strong nuclear force. Protons and neutrons are known collectively as nucleons. If individual nuclei have enough kinetic energy to overcome the repulsive Coulomb force of the protons, a single bound nucleus can be formed from the separate components. In environments in which the temperature is high enough, such as in the cores of stars, nuclei are able to fuse together (or burn) to produce all the known elements starting from the simplest atom, hydrogen. The progress made in understanding the physics of nucleosynthesis is described in the next section.

2. Nucleosynthesis

Each chemical element is labeled by a unique one or two letter abbreviation. For example, hydrogen is labeled H, helium He and carbon C. The number of protons plus neutrons in an atom's nucleus is known as the atomic number A. Because protons and neutrons have nearly the same mass and are much heavier than electrons, the atomic number of an atom determines its mass. Depending on how many neutrons are in the nucleus, atoms of a given element can have different masses; these are called isotopes of an element. The atomic number is used to identify the different isotopes. For example, the commonest isotope of carbon is ^{12}C which has 6 protons + 6 neutrons in the nucleus. ^{13}C and ^{14}C each have 6 protons but 7 and 8 neutrons respectively.

The lightest elements are hydrogen, helium and lithium (Li) containing 1, 2 and 3 protons respectively in their nuclei. Hydrogen has three isotopes: the commonest form contains one proton only (no neutrons) and is labeled H, deuterium or D contains one proton and one neutron, while tritium or T has one proton and two neutrons. Deuterium and tritium are sometimes written as ^{2}H and ^{3}H respectively. The two isotopes of helium are ^{3}He and ^{4}He.

One of the first attempts to develop a general theory of element synthesis was made by von Weizacker (1937). He suggested that all elements were built up from hydrogen in the deep interiors of stars by a series of neutron bombardments and decays. He suggested that these nuclear reactions were the source of stellar energy.

His work was extended by Bethe (1939) who presented the first detailed theory of thermonuclear energy generation in the Sun. Bethe showed how to make helium from hydrogen, but because there was no stable isotope containing five nucleons he could not produce any elements beyond ^{4}He. Bethe went so far as to say that "... under present conditions, no elements heavier than helium can be built up in the stars to any appreciable extent".

An alternative scenario for element synthesis was proposed by Gamow (1935). Building on Lemaître's ideas, he suggested that the elements were produced from a primordial neutron soup during a hot, dense phase in the early Universe, rather than in the interiors of stars. He used the word *ylem* (from a Greek term meaning 'first material') to describe this state of matter. When new nuclear data became available in 1946, Gamow gave the problem of element synthesis in a hot neutron soup to his student Alpher. Alpher set about calculating the abundances that could be expected from neutron capture in a hot Universe. The results were published in the *Physical Review* on 1st April 1948 (Alpher, Bethe & Gamow 1948). Because Bethe had not actually contributed to the work (Gamow's mischievous sense of humour led him to invite Bethe to join the author list), his name was supposed to be followed by "(in absentia)", but through some oversight this got erased in the final editing. This well-known publication is known as the $\alpha\beta\gamma$-paper.

The $\alpha\beta\gamma$-paper could account for the abundances of elements found in the Universe, attributing them to the period shortly after the Big Bang when nucleosynthesis occurred. Fermi and his student Turkevich redid the computations in the $\alpha\beta\gamma$-paper using exact nuclear data rather than the smoothed averages used by Alpher. They showed that even in the Big Bang scenario there was a problem producing elements beyond He (Fermi & Turkevich 1950). Another inaccuracy in the calculations of Gamow's group was that they assumed pure neutrons as an initial condition. Hayashi (1950) recognized that together with the neutrons, protons precipitated out of the plasma and played an important role in Big Bang nucleosynthesis (BBN). While the data included in the models has improved over the years, the framework for BBN calculations has not changed significantly since Hayashi's work.

A method for synthesizing elements heavier than helium was found by Salpeter (1952). His solution was the triple-alpha process, in which three helium nuclei, also known as alpha particles, combine to form carbon. This process requires three helium nuclei to come together at the same time, and can only occur when the density and temperature are high. Three-body collisions are rare events, and hence high densities are required to provide a reasonable chance for enough reactions to occur, and high temperatures are needed to give the nuclei enough energy to overcome the repulsive electrostatic forces between the nuclei. Although the temperature was high enough in the early Universe for this process to operate, the density was way too low and, therefore, essentially no C or any other metals (astronomers refer to all elements other than H and He as metals) were made during the BBN phase. The right conditions for the triple-alpha process to work exist in the centres of stars with masses above $\sim 2\,M_\odot$ ($M_\odot = 2 \times 10^{30}$ kg is the mass of the Sun) once all the H in the core has been used up. Once C has been made, the door is open for further heavy element synthesis via processes such as carbon burning, oxygen burning and neutron capture.

A landmark paper on the subject of nuclear synthesis in stars was published by Burbidge, Burbidge, Fowler & Hoyle (1957). The paper is now known as B^{2}FH from the initials of its authors. Working independently, Cameron (1957a,b) also laid down the fundamentals of nucleosynthesis. Unfortunately, Cameron was doing classified work at the time so the second of his papers was restricted for many years, and hence he did not share in much of the early recognition. Each of the papers proposed about ten processes by which elements were synthesized sequentially in the stars, resulting in a gradual conversion of hydrogen into heavier elements. In the B^{2}FH paper no allowance was made for synthesis in a hot Big Bang. During the late 1950s stellar synthesis rather than the Big Bang was the favoured mechanism for the origin of the elements.

In the steady state theory all elements other than hydrogen would need to be synthesized in stars. Ironically, Hoyle was one of the first to recognize that the observed abundance of helium presented a problem for the steady state theory (Hoyle & Tayler 1964). The measured ratio of the number of He atoms to the number of H atoms in a variety of astronomical objects was found to be of order ~ 0.1. This uniformly high abundance of He is difficult to reconcile with the small (~ 0.02) number of heavy atoms (metals) and the luminosities of the stars if stellar nucleosynthesis is responsible for the production of all elements heavier than hydrogen. Hoyle conceded that the only way to synthesize so much He required a far more dramatic mechanism than nuclear burning in the cores of stars. He and Fowler, together with graduate student Wagoner, set about calculating the abundances of all the light elements that would be formed in the Big Bang. Before their results were published, however, a far more important discovery was made that swayed opinion back in favour of the Big Bang.

3. The cosmic microwave background radiation (CMBR)

Two researchers working for the Bell laboratory, Penzias & Wilson, were trying to calibrate a horn antenna by pointing it at a patch of blank sky. They expected to find a region where they would get no signal which could be used to set a zero point, but regardless of the direction in which they pointed their feed, they always measured a weak signal corresponding to an object at a temperature of ~3 K. Penzias & Wilson (1965) published their results but offered no explanation of where this radiation in the microwave band came from.

Theoretical calculations by Alpher et al. (1948) predicted that a hot phase in the early Universe would leave behind a faint glow that would pervade space uniformly in all directions. Dicke and co-workers at Princeton University had come to the same conclusion independently and had built equipment to search for this cosmic radiation. At that stage the age of the Universe was still very uncertain. Dicke and co-workers where searching for background radiation at a temperature of ~30 K and, therefore, were using higher frequencies than those used by Penzias & Wilson. The CMBR discovered by Penzias & Wilson was recognized by Dicke et al. (1965) as being the remnant radiation left over from a hot fireball stage in the early Universe.

The discovery of the CMBR sounded the death knell for the steady state theory. This radiation cannot be accounted for in the steady state theory, but is a consequence of a Universe that was hot in its early phase. Once its existence had been established the next question was what elements were produced by this event, and what were their abundances. Peebles (1966) was able to estimate what the relative abundance of primordial D and He would be, while Wagoner, Fowler & Hoyle (1967) did similar calculations and showed that the production of deuterium was unique to the Big Bang.

4. Big Bang nucleosynthesis

In the Big Bang scenario, the Universe was created at a time $t = 0$ in a hot, dense state, and started expanding. After ~10^{-43} s the plasma consisted of elementary particles such as quarks, gluons, neutrinos, electrons, positrons and photons. Because of the high temperatures, particles were moving relativistically and hence collisions occurred frequently. As a result, equilibrium was established and maintained; any deviations from equilibrium were erased quickly. Depending on the energies associated with different reactions, the various species of particle maintained equilibrium ratios until the temperature cooled below some critical value. For example, the neutron-to-proton ratio was maintained at its equilibrium value

$$n/p = e^{-\Delta m/T}$$

where Δm is the difference in mass between the proton and the heavier neutron, by reactions such as:

$$p + e^- \leftrightarrow n + \nu_e$$
$$n + e^+ \leftrightarrow p + \bar{\nu}_e \text{ and}$$
$$n \rightarrow p + e^- + \bar{\nu}_e$$

By $t = 1$ s the temperature was $T \approx 10^{10}$ K. Reactions converting neutrons to protons still occurred but the reactions involving neutrinos (ν_e and $\bar{\nu}_e$) were too slow to maintain equilibrium. The neutron-to-proton ratio effectively "froze out" of the gas.

By $t = 10$ s the temperature had reached $T \approx 3 \times 10^9$ K and occasional nuclear reactions between protons and neutrons were occurring to produce deuterons by the reaction $n + p \rightleftharpoons D + \gamma$. The symbol $\rightleftharpoons$ indicates that this is a reversible reaction; the deuteron can be broken apart if it is hit by a γ-ray in a process called photodissociation. For photodisintegration of a deuteron to occur, the incident γ-ray photon must have an energy that exceeds 2.23 MeV. Although at $t = 10$ s the average energy of the photons was less than this, because the nucleons were merely a trace contaminant in a bath of radiation ($n_N / n_\gamma = \eta < 10^{-9}$), there was a sufficient number of photons with energies $E > 2.23$ MeV that any deuterons formed were rapidly dissociated. This kept the D abundance very small and imposed a bottleneck on any further nucleosynthesis. Thus, roughly 10 s into the evolution of the Universe, no significant nucleosynthesis had yet occurred.

Note that in the early Universe there were almost as many free neutrons present as there were protons. Deuterium was able to form when a proton and neutron collided with one another. In the cores of main-sequence stars, very few isolated neutrons are present and so D is formed by a different reaction to that operating in the BBN phase. The appropriate reaction is $H + H \rightarrow D + e^+ + \nu$, that is two protons collide to produce a deuteron, a positron e^+ and a neutrino ν. The presence of a neutrino in this reaction indicates that the reaction involves the weak nuclear force and hence this is a process with a small interaction probability. This reaction did not contribute much to the production of D in the early Universe.

As the Universe expanded and cooled further, fewer and fewer photons were capable of dissociating the deuterons. Hence, the D abundance increased and by $t = 100$ s and $T \approx 10^9$ K the deuterium bottleneck had been overcome. Big Bang nucleosynthesis (BBN) began in earnest. Tritium and ^{3}He were formed, and these in turn fused into ^{4}He. Some trace amounts of ^{7}Li and ^{7}Be were also produced. After 1000 s had elapsed, the temperature dropped below $T \approx 10^9$ K and nucleosynthesis ceased. Further nucleosynthesis had to wait until stars condensed, and the temperatures and pressures in their cores reached sufficiently high values that nuclei could be forced to merge. Computed abundances of the light nuclei during the first few minutes of the Universe are illustrated in Fig. 1 as a function of time and temperature.

Because the baryon (nucleon) density during the BBN phase was relatively low (about 1% the density of terrestrial air), only reactions involving two-particle collisions occur. Combining the most abundant nuclei, protons and ^{4}He, via two body interactions always leads to unstable mass-5 nuclei. Even when ^{4}He combines with rarer nuclei like ^{3}H or ^{3}He, mass-7 nuclei are formed which, when hit by a proton, yield mass-8 nuclei. These nuclei are unstable and decay to ^{7}Li very rapidly. Only when gravity compresses the matter to sufficient densities, can three-body collisions, such as the triple-alpha process ($3 \times {}^4\mathrm{He} \rightarrow {}^{12}\mathrm{C}$), occur to make heavy elements. Thus, BBN makes ^{4}He with traces of ^{2}D, ^{3}He, ^{7}Li and ^{7}Be.

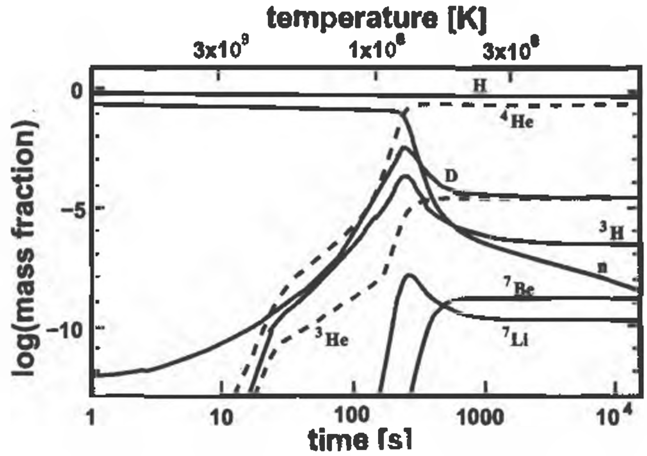

Figure 1. Theoretical calculation of the light element abundances as a function of time in the standard BBN model.

In standard BBN the abundances of the light elements D, ^{3}He, ^{4}He, and ^{7}Li depend on the number of relativistic neutrino species N_ν, the neutron lifetime τ_n, and the baryon-to-photon number ratio $\eta \equiv n_b / n_\gamma$ Experimental data give $N_\nu = 2.992 \pm 0.016$ (LEP Working Group 1995) and $\tau_n = 887 \pm 2$ s (Review of Particle Properties 1994). Thus, only $\eta \equiv n_b / n_\gamma$ is left to be determined experimentally. The largest contribution to the photon density, n_γ, comes from the CMBR, and hence n_γ is directly related to the temperature of the microwave background. This has been measured accurately using the COBE satellite at $T = 2.726 \pm 0.005$ K (Mather et al. 1994), which then gives $n_\gamma = 411$ cm^{-3}. Measurements of the primordial light element abundances therefore allow a determination of the baryon density of the Universe. This is done by comparing predicted abundances of D, ^{3}He, ^{4}He, and ^{7}Li from a grid of BBN models for varying η with the corresponding primordial values inferred from measured abundances.

Several model calculations have been performed in which some of the constraints discussed above have been removed. For example, inhomogeneous models have been constructed which have dense proton rich zones co-existing with low density neutron rich ones. Dense zones would have produced low D, high ^{4}He abundances, while dilute ones would have led to the reverse. Anisotropic expansion has been considered, as well as alternative theories of gravity. At present there is no compelling evidence to suggest that any of these alternatives need to be considered. Refinements in the measurements might help to constrain these alternative theories.

5. Evolution of stars and galaxies

The determination of primordial abundances from astronomical observations is complicated by the fact that, in the course of their evolution, stars mimic some of the processes that occurred during the BBN phase, thereby polluting the primordial material with products of stellar nucleosynthesis. D and ^{3}He are consumed by stars during their life, and hence the amounts of these elements are reduced as the Universe ages. During their main-sequence phase, stars burn H to produce ^{4}He. When stars die, they return this processed material to the interstellar medium (ISM), thereby increasing the abundance of ^{4}He. If primordial abundances are to be inferred from present day observations, then it is important to understand what processing of material has occurred since BBN. In this section the effects of stellar and galactic evolution are discussed to indicate how primordial abundances are affected.

Stars form out of giant clouds of dust and gas that collapse due to gravitational instabilities within the gas. The initial mass of a star determines many of its vital statistics such as brightness, temperature, and lifetime. The reason for this is that the mass determines its core temperature, which in turn determines the rate at which nuclear reactions occur. The rate at which energy is produced by nuclear fusion reactions determines the brightness and surface temperature of the star. The faster the reactions, the sooner the fuel will be used up and hence the shorter the star's lifetime. Whereas low mass stars are only able to burn H to He, the high core temperatures found in massive stars are able to synthesize metals.

The processes by which stellar nucleosynthesis occurs were described in the papers by B^{2}FH and Cameron in 1957. A condensation of gas with a mass $M > 0.08\ M_\odot$ will attain a core temperature exceeding 10^7 K. Above this temperature the kinetic energy of protons in the gas are sufficient to overcome the repulsive Coulomb force of the positive charges and H nuclei can be fused to form He. This initial stage of any star's life, when H is being burned to produce He in its core, will last for 80–90% of its lifetime. Stars in this phase of life are very common and are known as main-sequence stars. Once the H in the core has been used up, the core contracts, and heats up. If the star has a large enough mass that the core temperature exceeds 10^8 K, helium burning (via the triple-alpha process) will occur, producing C and O in the core. This is the main source of energy for stars in their red-giant phase. During the main-sequence and red-giant phases of stellar evolution, stars return some of their material to the ISM via stellar winds. This matter may contain processed material that has been dredged up from the interior of the star by convection. If $M_* > 8\ M_\odot$ the temperature in the core exceeds 10^9 K, and carbon/oxygen burning will occur, producing elements such as neon (Ne), sodium (Na), magnesium (Mg), silicon (Si) up to the formation of an iron-nickel (Fe-Ni) core. Collapse of the core follows, resulting in a supernova explosion. A large fraction of the mass of the star is ejected, leaving behind a neutron star or black hole. In stellar cores, element synthesis stops at Fe; all elements in the periodic table heavier than Fe are produced by rapid neutron capture processes as a result of a supernova explosion. The bulk of the observed heavy metals in the Universe have been produced by supernova explosions of stars with $M > 10\ M_\odot$.

A condensation of gas with a mass $M < 0.08\ M_\odot$ will never attain a temperature in its core sufficient to ignite H. This type of object is known as a brown dwarf. Brown dwarfs lock up material and do not participate in galactic enrichment. Even stars with $M < 0.5\ M_\odot$ have lifetimes that are much longer than the Hubble time (the age of the Universe, presuming it has always expanded at the currently observed rate) and, hence, they too do not participate in galactic evolution. Stars with $M > 20\ M_\odot$, on the other hand, live for less than 20 million yrs, and end their lives spectacularly in supernova explosions that return to the ISM large amounts of processed material. This enriched matter might fall onto other stars polluting the outer layers, or it might be incorporated into clouds of dust and gas out of which new stars will condense. Attempts to model this cycling of material are complicated by the large number of free parameters contained in galactic chemical evolution models.

Although the finer details of galactic chemical evolution are difficult to reconcile with observations, it has been recognized for many years that stars in the spiral arms of galaxies are richer in heavy elements (metals) than stars in galactic bulges or in elliptical galaxies. The stars in the spiral arms are called Population I stars, those in galactic bulges and elliptical galaxies Population II stars. The high metallicity of the Population I stars suggests they were formed from material that has been cycled through stars more times than is the case for the Population II stars, and hence Pop I is believed to be younger than Pop II stars. An even older generation of stars is thought to have existed before the Population II stars formed. These Population III stars, as they are called, would have formed out of primordial material and hence would contain no metals, only H and He. Objects with low metallicity should have the least amount of contamination and hence are the best candidates to use for primordial abundance determinations.

6. Determining abundances

When determining abundances it is usual to measure the ratio of the number of atoms of a particular element or isotope relative to the number of H atoms. The abundance of lithium relative to hydrogen is written N(Li)/N(H) or simply Li/H. For He this quantity is labeled by y, i.e.

$$y = \frac{N(^4\mathrm{He})}{N(\mathrm{H})}.$$

In ionized gases (H II regions) the abundance of ionized helium, He^+, relative to ionized hydrogen, H^+, is measured which is denoted by

$$y^+ = \frac{N(^4\mathrm{He}^+)}{N(\mathrm{H}^+)}.$$

Conventional practice quotes the He abundance in terms of its mass fraction rather than its relative abundance by number. The symbol used for the mass fraction of He is Y, and because ^{4}He is about four times heavier than H and together with H makes up the bulk of the mass in the Universe, it follows that

$$Y = \frac{4y}{1+4y}.$$

All elements other than H or He are usually lumped together and referred to as "metals" by astronomers. In particular, oxygen, the next most abundant element in the Universe after He, usually produces the most prominent lines in nebular gases and hence is often used as a measure of the metallicity in H II regions. The mass fraction of metals is denoted by Z.

To determine primordial abundances of the light elements, a number of steps are involved. To start with a measurement has to be made. Because the signals are often weak, the observations are difficult and require careful extraction of the signal from the noise.

Once a measurement has been made, the next step is to convert it into an abundance. This is usually a non-trivial exercise. For example, when using the emission lines from a cloud of ionized gas to determine abundances, the intensity of the lines provides a measure of the number of ionized atoms in the gas. For example, the lines of He I provide a measure y^+. There could also be some neutral helium, He^0, in the gas, which is undetectable, and some He^{++}. To determine an "observed" mass fraction of He, Y_o, some way of estimating the level of ionization is needed to correct the measured $N(He^+)$ for the amount of He in unobserved stages of ionization.

In the final stage, a primordial value Y_p has to be inferred from the observed abundance Y_o. To do this some estimate of the amount of processing that the material has been subjected to has to be made. The metallicity of an astronomical object provides a measure of how much processing has occurred, but it is also necessary to know what effect this processing has had on the primordial abundances. Sources and sinks of isotopes of any element need to be accounted for, which requires a theoretical understanding of production and destruction mechanisms. For each element different mechanisms operate and so they need to be considered separately. In the next section the processes altering the primordial abundances of the light elements are discussed, and a brief summary of observed values are presented.

7. Abundances

For all the elements, solar system abundances, particularly terrestrial values, are the easiest to measure, but are representative of the space-time volume 4.65 Gyr ago and 6-10 kpc from the Galactic centre. Extrapolating these values back to the first few minutes after the Big Bang cannot be done with any confidence however, but the local values do provide a convenient limit to any other astronomical estimations. To appreciate how observed abundances can be used to constrain Big Bang models, the results of theoretical calculations are presented before going on to look at the measured values.

7.1 Theoretical Models

Calculated primordial abundances using a standard BBN model are shown in Fig. 2. Relative to the number of H atoms, ^4He contributes only 8–9% of the atoms in the Universe, but this translates into a mass, $Y \approx 0.25$, or 25% by mass of the baryonic content of the Universe. Although the number abundance is low, after H, helium is the next most abundant element in the Universe. A consequence of the large abundance of ^4He is that it can be observed throughout the cosmos in a variety of objects. Furthermore, its abundance can be determined to a higher accuracy than for other light elements. There is constancy in Y at the level of ±20% in all objects, indicating a universal origin (synthesis during the Big Bang) for this isotope.

Of the light elements, D is the next most abundant species but there are about 10^4-10^5 H atoms for ever D atom. The steep slope of the D/H curve shown in Fig. 2 indicates that this ratio is a sensitive indicator of the baryon density ($\eta = n_b/n_\gamma \propto \Omega_b h^2$) in the era of BBN. Because of its low abundance, however, observations of D are difficult and subject to large errors. This situation is even more pronounced for ^{3}He. Because there are mechanisms producing and destroying ^{3}He in stars, determinations of its primordial abundance are model dependent and not very reliable.

The characteristic valley shape of the ^{7}Li/H curve shown in Fig. 2 is a consequence of two competing production paths. Because of some uncertainties in several key reaction rates for the destruction of ^{7}Li, predicted abundances of this isotope are uncertain by a factor of ~2 – 3. The ratio Li/H ~ 10^{-10} – 10^{-9} means that Li measurements are difficult to make and subject to large uncertainties.

The figure also indicates the present limits for the abundances as determined from observations. The light element abundances are discussed in more detail in the next sections.

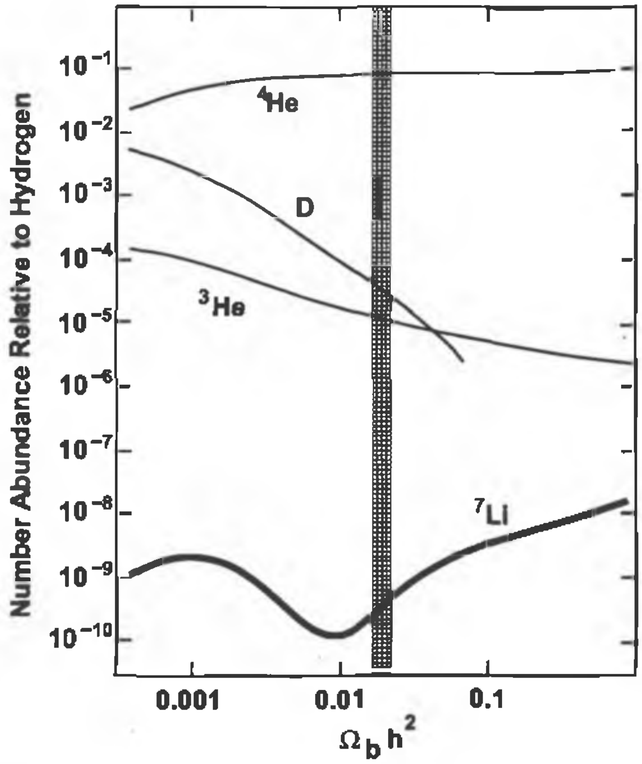

Figure 2. Theoretical calculations of the light element abundances from standard BBN models. The parameter $h = H_0/100$ where H_0 is the Hubble constant, and Ω_b is the density of baryons. The baryon-to-photon number ratio η is related to $\Omega_b h^2$.

7.2 Deuterium

Supernova shock waves and spallation (the breakup of a bombarded nucleus into several parts), amongst others, have been suggested as mechanisms capable of producing deuterium. Epstein, Latimer & Schramm (1976) considered the various alternatives and concluded that post Big Bang deuterium production requires extremely violent and exotic events, all of which are unlikely to produce any significant quantities of D. Uniquely among the light isotopes, D appears to have no sources of formation other than BBN.

Deuterium is produced as the first step in the PP I chain of H burning, but D is a fragile isotope and is easily destroyed in stars. If temperatures exceed $T \sim 6 \times 10^5$ K it is burned to ^{3}He. Convection occurs in pre-main sequence collapse and sufficiently high temperatures are reached to convert any D in the material to ^{3}He. To date, no D has been detected in stellar atmospheres, and hence it appears that any gas that has been cycled through stars converts all D to ^{3}He. Because D is readily destroyed in stars, observed abundances can only provide a lower limit on the primordial D abundance.

Until a few decades ago all that was known about the D abundance was that D/H on Earth was ~ 1/3000. This value appears to be enhanced by large amounts compared to other determinations within the solar system. Measurements of D in the atmospheres of the giant planets yield a pre-solar system abundance of D/H ~ 10^{-5}.

Although there exist extensive observations of deuterated molecules (DCO^+, DCN, DNC, HDO, etc.) in the ISM, virtually no light has been shed on the primordial abundance of D. The reason is that chemical and physical fractionation effects have led to enormous enhancements of D in various molecules. The observations of these molecules are of more relevance to interstellar chemistry than to the abundance determination of primordial deuterium.

The 1 420 MHz ($\lambda = 21$ cm) line of H is due to a hyperfine spin-flip transition. It has provided lots of useful information about the Milky Way and other galaxies. Like H, deuterium has a hyperfine spin-flip transition that produces radiation at 327 MHz ($\lambda = 92$ cm). After many searches in the cosmos for this emission, however, it has evaded detection thus far.

The first detection of pure D in the ISM came from UV observations made using the *Copernicus* satellite. Rogerson & York (1973) detected Lyman absorption lines of D towards the star Beta Centauri, from which they estimated a D/H value of 1 : 70 000. Subsequently, interstellar D has been found in the wings of the Lyman lines of H I along the lines of sight to several nearby cool stars and to more distant O and B stars.

The metal abundance in the intergalactic medium (IGM) is low, indicating that very little stellar processing has occurred in the gas. Hence, IGM absorption lines seen towards distant quasars should provide a good estimate of the primordial D abundance. Measurements of D in high redshift clouds have produced conflicting results, however. Kirkman et al. (2003) compiled a list of several such measurements and reported a value of $(D/H)_{\mathrm{prim}} = 2.78^{+0.44}_{-0.38} \times 10^{-5}$, although there is significant scatter in the individual measurements.

Sembach et al. (2004) found $D/H = (2.2 \pm 0.7) \times 10^{-5}$ in Complex C, a high-velocity cloud falling onto the Milky Way. It has low metallicity but presumably has experienced more stellar processing than most IGM material. As a result of multiple phases of stellar recycling, the D/H value within the local disk of the Milky Way is significantly lower than in the IGM or Complex C (Wood et al. 2004).

7.3 ^{3}He determinations

When stars form out of molecular clouds, any deuterium present in the pre-stellar cloud will be burned to ^{3}He during the collapse of the cloud. Hence, the initial amount of ^{3}He will be augmented by the D incorporated in the star. ^{3}He is burned away in the interiors of stars, but some survives in the outer layers of stellar atmospheres. ^{3}He is also a product of incomplete hydrogen burning in the cooler outer regions of stars with $M < 2.2\, M_{\odot}$ (Iben 1967). Because ^{3}He is both destroyed and produced by stellar processing, the chemical galactic evolution of

^{3}He is more complicated than that of D.

The solar wind has implanted ^{3}He into gas-rich meteorites and lunar soil and breccias. However, the solar wind abundance of ^{3}He is representative of the pre-solar abundance of D + ^{3}He. The solar wind may have been contaminated by additional ^{3}He dredged up from the interior, but it is difficult to estimate the importance of this effect. The solar system abundance determination of ^{3}He provides an upper limit to the primordial abundances of $(D + {}^3He)/H = 3.6 \times 10^{-5}$.

Singly ionized ^{3}He has a spin-flip transition analogous to the 21 cm line of H and 92 cm line of D. Although the 8.7 GHz ($\lambda = 3.46$ cm) emission from this transition has been found in some Galactic H II regions (Rood, Bania & Wilson 1984), it is weak, and the small line-to-continuum ratio makes it difficult to determine the ^{3}He abundance. Furthermore, it is far from trivial to derive a meaningful $^3He/H$ ratio from $^3He^+$. The mean value determined by this method is consistent with other determinations but there is a large scatter in the results.

Gloeckler & Geiss (1996) measured the $^3He/^4He$ abundance ratio in the local interstellar cloud from an isotopic analysis of helium atoms that have entered the solar system from the surrounding ISM. The measurements were made with the Solar Wind Ion Composition Spectrometer (SWICS) aboard the *Ulysses* spacecraft during a 40 month period. They found an $^3He/^4He$ number density ratio in the local interstellar gas of 2.2×10^{-4} which agrees with the mean value of $^3He/H$ (with $^4He/H \approx 0.1$) found from the hyperfine line measurements of $^3He^+$ in H II regions (Rood et al. 1984).

Rood et al. (1998) suggested $(^3He/H)_p = 1.5^{+1.0}_{-0.5} \times 10^{-5}$ as a reasonable estimate for the primordial ^{3}He isotope. More recently, Bania, Rood & Balser (2002) report a limit on the primordial $^3He/H$ ratio from their detailed long term study of 60 Galactic H II regions and six planetary nebulae. In 17 Galactic H II regions for which the ionization corrections were relatively simple, they find a mean $^3He/H = (1.9 \pm 0.6) \times 10^{-5}$, which will be an upper limit on the primordial ratio if stars have not on average destroyed ^{3}He. They propose that the best value for the upper limit on the primordial $^3He/H$ is the value they measured for one H II region that has the lowest metal abundance in their sample, the third lowest $^3He/H$ ratio, excellent data, and a small ionization correction of 22%. They then quote $^3He/H < (1.1 \pm 0.2) \times 10^{-5}$, which is consistent with the value predicted by D/H and standard BBN. Even this latter, more precise, determination of $(^3He/H)_p$ does not constrain the value of η significantly because of the weak dependence of this ratio on the baryon density.

7.4 ^{7}Li determinations

A large fraction of Li present today is believed to have been produced by nucleosynthesis in the early Universe. However, there appears to be a source besides the Big Bang that produces Li and the next two light elements beryllium (Be) and boron (B). These elements are fragile and rapidly disintegrate in the hot, dense interior of most stars. Therefore, it is surprising to find that their abundance in the spectrum of galactic cosmic rays is enhanced by as much as a million compared to the ISM value. The interaction of cosmic rays with the ISM has emerged as the most plausible of several candidates to account for the synthesis of Li, Be, and B. Light cosmic rays (protons and alpha particles) can interact with heavy nuclei (C, N, O) in the ISM, or heavy nuclei can interact with light nuclei. Cosmic ray models lead to predictions of the absolute abundances of ^{6}Li, ^{9}Be, ^{10}B and ^{11}B that are quantitatively close to the observed abundances. However, the ratio $^{10}\mathrm{B}/^{11}\mathrm{B}$ is known to a much higher degree of accuracy than the absolute abundances, and is not well accounted for by the models. The predicted value is 2.5, the measured value is 4.05.

The maximum Li/H value in Population I stars from the time of the origin of the solar system (as measured in meteorites) to the present (as found in the youngest stars and the interstellar gas) seems to have remained remarkably constant at a value of $\mathrm{Li/H} = 10^{-9}$. The nett effect of Li destruction and Li production in the Galactic disc on the overall Li abundance appears to have been small. However, the measured Li/H value is a combination of BBN and Galactic cosmic-ray produced ^{6}Li and ^{7}Li.

Many high accuracy Li/H abundance determinations have been made for the atmospheres of metal poor halo stars. There is very little scatter, and hence the abundance ratio in these stars is well determined at $(1-2)\times 10^{-10}$ (Ryan et al. 2000 and references therein). No Be or ^{6}Li has been detected in these stars, indicating that they have not been contaminated by galactic cosmic-ray spallation products. The measured Population II Li/H value probably represents a lower limit for ^{7}Li. The actual value of the primordial ^{7}Li abundance is still controversial.

7.5 ^{4}He determinations

^{4}He is the end product of H burning in main-sequence stars and only a small fraction of this extra component is converted to heavier elements via the triple-alpha process. Therefore, the overall abundance of ^{4}He increases with time. Abundance determinations and extrapolation to Y_{p} in all objects is model-dependent to various degrees, as discussed below.

In the solar system, measurements of ^{4}He in meteorites, the giant planets and in the Sun set an upper limit on the proto-solar value of $Y = 0.26 \pm 0.03$. Because the energy gap between the two lowest levels of He is larger than in H, absorption lines of He I and He II can only be seen in the hottest stars (type O and B). These are young stars that have only recently formed, and hence their He content has been modified by previous generations of stars. ^{4}He abundances have been determined in H II regions in the Milky Way, but their metallicity is high and extrapolating back to a primordial value has lots of uncertainties. These studies have provided much useful information on the physics of nebulae, though. Likewise, He/H has been measured in planetary nebulae (PNe) but not much is learned about the primordial He abundance. There will be contamination of the primordial material due to processing by the central star of the given PN, and due to the general Galactic chemical enrichment. Although there is good agreement among the various planetaries, many model-dependent assumptions about galactic and stellar evolution must be made to estimate a primordial abundance.

The stars in many globular clusters are known to be old and metal-poor. Furthermore, there are several features of the colour-magnitude diagrams that are sensitive to the He abundance, such as the width of the instability strip, the luminosity of the main-sequence turn-off point, and the luminosity of the horizontal branch. However, there are large uncertainties and systematic errors in estimating the effects due to age, metallicity and internal mixing in the stars of globular clusters, and in the fitting of cluster diagrams to theoretical curves. An application of the method to much more complete data sets and with updated stellar models indicates $Y_p \equiv 0.243 - 0.244$ (Cassisi, Salaris & Irwin 2003).

Using the Hubble Space Telescope, Jakobsen et al. (1994) detected strong absorption arising from singly ionized helium (He II) along the line of sight to quasar Q0302–003 which has a redshift $z = 3.286$. The strength of the absorption suggests that it may arise in a diffuse ionized intergalactic medium. Using the 10-metre Keck telescope and a higher resolution spectrograph, Songaila, Hu & Cowie (1995) also looked at quasar Q0302–003. They detected a population of weak clouds with column density down to 2×10^{12} cm^{-2}, and showed that absorption in these clouds can account for the strong He II absorption without having to invoke a diffuse intergalactic medium. The detection confirms that substantial amounts of helium existed in the early Universe, as predicted by BBN, and that the Lyman forest clouds (and intergalactic medium, depending on whose interpretation is used) are indeed highly ionized. Unfortunately, this high level of ionization also means that the observable species of H I, He I and He II are mere trace constituents of the intergalactic gas, and that most of the mass is in the form of unobservable H II and He III. Due to the huge uncertainties associated with the large ionization corrections, the prospects for determining an accurate value for the primordial helium abundance are rather poor.

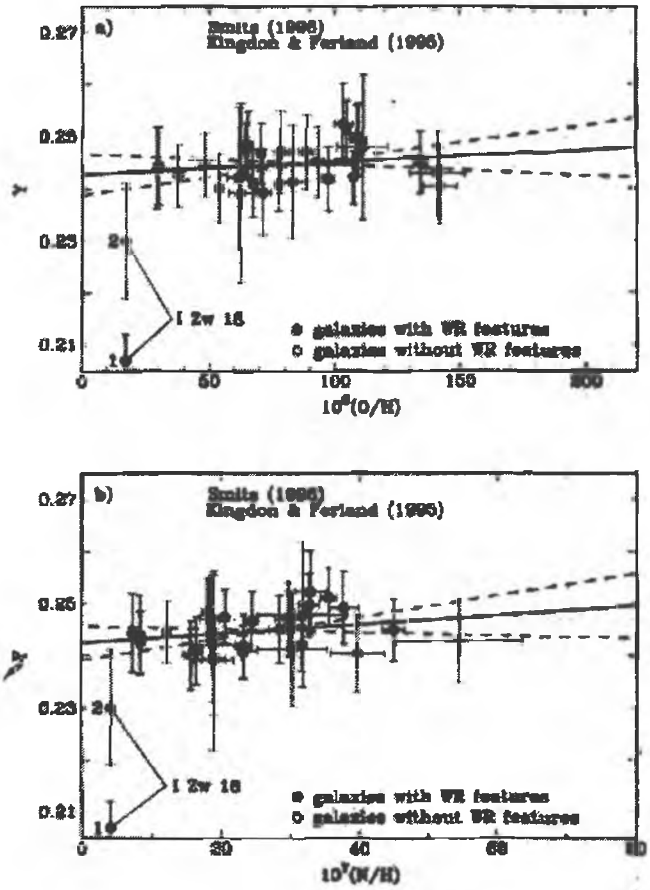

Figure 3. Plots of He versus O and N abundances to determine Y_p (from Izotov et al. 1997).

Currently, the most accurate determinations of Y_p are being made on objects known as blue compact galaxies (BCGs) containing giant extragalactic H II regions. For many years only one such object was known – I Zw 18 – but surveys have led to the discovery of many more galaxies of this class.

BCGs are objects with $M_B \geq -18$ presently undergoing intense bursts of star formation that give birth to ~ $10^3 - 10^4$ massive stars in a compact region. The optical spectra of BCGs have strong, narrow emission lines superposed on a stellar continuum rising toward the blue. The radiation from the hot young stars gives rise to the blue colour of the continuum, and the supergiant H II regions in which the stars have formed produce the nebular emission line spectrum. The metallicities of BCGs fall in the range $Z_\odot/50 \leq Z \leq Z_\odot/3$ where $Z_\odot$ is the solar metallicity abundance. These low metallicities imply that their gas has not been processed through stars many times, and hence that they are relatively young from the point of view of chemical evolution. The gas in the H II regions has not been polluted by non-primordial helium synthesized in stars. The observed [O III]/[O II] intensity ratios are fairly large, implying that the gas is in a state of high excitation. The λ4 868 line of He II is seen in the spectra, providing a measure of the He^{++} abundance, while the high excitation means that corrections for H^0 and He^0, the unseen neutral components mixed with the ionized gas, are small.

The y^+ abundance is determined by comparing the measured fluxes of He I emission lines with theoretical values for the temperature and density derived from the forbidden line fluxes in the gas. The amount of y^{++} is determined from the intensity of the He II λ4 868 line. The observed mass fraction of helium, Y_0, thus obtained has to be converted to a primordial value Y_p. A standard approach, based on an idea of Peimbert & Torres-Peimbert (1974), is to assume that the amount of helium produced by stellar processing is linearly proportional to the amount of O or N in the gas. An H II region with $Z = 0$ would have a Y-value of Y_p, while in H II regions with $Z > 0$ the Y-value will have increased by an amount ΔY. The chemical evolution of a galaxy can be characterized by a relation of the form $\Delta Y = \alpha \Delta Z$, where α depends on time and position. Hence, by plotting the He abundance against the O or N abundance for a number of BCGs, and extrapolating to zero metal abundance, a value for the primordial abundance of He can be obtained.

An example of this procedure, using a linear relation of the form

$$Y = Y_p + \left(\frac{\mathrm{O}}{\mathrm{H}}\right) \frac{dY}{d(\mathrm{O/H})}$$

and similarly for N, is shown in Fig. 3. This diagram is taken from the work of Izotov et al. (1997) who used the theoretical He calculations of Smits (1996) to determine the helium abundances.

The best current estimate of the primordial He abundance determined by Izotov & Thuan (2004) depends to some extent on some assumptions made about the data. Values lie between $Y_p = 0.2421 \pm 0.0021$ and $Y_p = 0.2444 \pm 0.0020$. These give baryonic mass fractions of $\Omega_b h^2 = 0.012^{0.003}_{0.002}$ and $0.015^{0.003}_{0.002}$, which are lower than values derived from the D abundance and fluctuations in the CMBR. The source of the discrepancy is under investigation.

8. New methods

On large scales the Universe is highly uniform, but on small scales this is not the case. Anisotropies in the intensity of the CMBR have been observed at a level of a few parts in 100 000 over a range of angular scales. These tiny variations in intensity represent fluctuations in the temperature of the CMBR, which in turn indicate the presence of density fluctuations in the distribution of cosmic matter. A statistically useful way of plotting the fluctuations is in the form of an angular power spectrum.

The predicted shape of the angular power spectrum depends on a number of cosmological parameters, including the Hubble constant H_0 and the baryon density Ω_b. By comparing a large number of computed power spectra with observational data, a 'best-fit' spectrum can be identified. This enables accurate testing of cosmological models from a totally independent standpoint. After a year of operation, data obtained with the *Wilkinson Microwave Anisotropy Probe* (WMAP) have provided the most accurate angular power spectrum to date. The data are regarded by many cosmologists to represent the best numerical characterization of the Universe currently available.

Using only WMAP data, the best fit values of the Hubble constant and baryon density are $H_0 = 72 \pm 5$ km s^{-1} Mpc^{-1} and $\Omega_b h^2 = 0.024 \pm 0.001$, giving a baryon-to-photon ratio of $\eta = 6.1^{+0.04}_{-0.03} \times 10^{10}$ (Spergel et al. 2003). Fitting the model parameters to a combination of WMAP data and other finer scale cosmic microwave background experiments such as the Arcminute Cosmology Bolometer Array Receiver (ACBAR), the Cosmic Background Imager (CBI) and the 2dFGRS measurements amongst others, gives $h = 0.71^{+0.04}_{-0.03}$ and $\Omega_b h^2 = 0.0224 \pm 0.0009$. These parameters imply that the age of the Universe is 13.7 ± 0.2 GYr. The D/H and ^{3}He/H primordial abundance determinations are consistent with the WMAP data (within the observational errors), but the measured ^{4}He/H and ^{7}Li/H values are too low. Although there are still significant uncertainties in some of the measured values of the light element abundances, the overall agreement from several independent lines of investigation is good. The term **precision cosmology** was coined in 1996 by the cosmologist Michael Turner. It expresses the direction in which cosmology is headed. Gone are the days when cosmologists were free to speculate about the origin and evolution of the Universe without any observational data to fetter them. We can only wait to see what mysteries the next generation of telescopes such as Planck, due for launch in 2007, will reveal.

References

Alpher R.H., Bethe H. & Gamow G. (1948) Phys. Rev., 73, 803. (αβγ-paper)
Alpher R.H. & Herman R.C. (1950) Rev. Mod. Phys., 22, 153.
Bania T. M., Rood R.T. & Balser D. S. (2002) Nature, 415, 54-57.
Bethe H. (1939) Phys. Rev., 55, 434.
Bondi H. & Gold T. (1948) MNRAS, 108, 252.
Burbidge E.M., Burbidge G.R., Fowler W.A. & Hoyle F. (1957) Rev. Mod. Phys., 29, 547. (B^{2}FH)
Cameron A.G.W. (1957a) PASP, 69, 201.
Cameron A.G.W. (1957b) *Chalk River Report* CRL-41.

Cassisi S., Salaris M. & Irwin S.W. (2003) ApJ, 588, 862.

Dicke R.H., Peebles P.J.E., Roll P.G. & Wilkinson D.T. (1965) ApJ, 142, 414.

Epstein R.I., Latimer J.M. & Schramm D.N. (1976) Nature, 263, 198.

Fermi E. & Turkevich A. (1950) Unpublished work described in Alpher & Herman (1950).

Gamow G. (1935) Ohio J. Sci., 35, 406.

Gloeckler G. & Geiss J. (1996) Nature, 381, 210.

Hayashi D. (1950) Prog. Theor. Phys., 5, 224.

Hoyle F. (1948) MNRAS, 108, 372.

Hoyle F. & Tayler R.J. (1964) Nature, 203, 1108.

Hubble, E. P. (1929) Proc. Nat. Acad. Sci USA, 15(3), 168-173.

Iben I. (1967) ApJ, 147, 624.

Izotov Y.I., Thuan T.X. & Lipovetsky V.A. (1997) ApJS, 108, 1.

Izotov Y.I. & Thuan T.X. (2004) ApJ, 602, 200.

Jakobsen P., Boksenberg A., Deharveng J.M., Greenfield P., Jedrzejewski R. & Paresce F. (1994) Nature, 370, 35.

Kirkman D.,Tytler D., Suzuki N., O'Meara J.M. & Lubin D. (2003) ApJS, 149, 1.

Lemaître G. (1931) MNRAS, 91, 483.

LEP Working Group (1995) Preprint (CERN PPE/95-172).

Mather J.C. et al. (1994) ApJ, 420, 439.

Peebles P.J.E. (1966) ApJ, 146, 542.

Peimbert M. & Torres-Peimbert S. (1974) ApJ, 193, 327.

Penzias A.A. & Wilson R.W. (1965) ApJ, 142, 419.

Review of Particle Properties (1994) Phys. Rev. D, 50, 1173.

Rogerson J.D. & York D.G. (1973) ApJ Lett., 186, L95.

Rood R.T., Bania T.M. & Wilson T.L. (1984), ApJ, 280, 629.

Rood R.T., Bania T.M., Balser D.S. & Wilson T.L. (1998) Space Sci. Rev., 84, 185.

Ryan S., Beers T.C., Olive K.A., Fields B.D. & Norris J.E. (2000) ApJ, 530, L57.

Salpeter E. (1952) ApJ, 115, 326.

Sembach K.R., et al. (2004), ApJS, 150, 387.

Smits D.P. (1996) MNRAS, 278, 683.

Songaila A., Hu E.M. & Cowie L.L. (1995) Nature, 375, 124.

Spergel, D. N. et al. (2003) ApJS, 148, 175-194.

von Weizacker C. (1937) Phys. Zs., 38, 176.

Wagoner R.V., Fowler W.A. & Hoyle F. (1967) ApJ, 148, 3.

Wood B.E., Linsky J.L., Gèrrard G., Williger G.M., Moos H.W. & Blair W.P. (2004) ApJ, 609, 838.